
The role of outflows driven by young stars and brown dwarfs investigated via forefront observations

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“Even through your hardest days, remember we are all made of stardust.”

- Carl Sagan

Abstract

The role of outflows driven by young stars and brown dwarfs investigated via forefront observations

by Noah OTTEN

Stars like our Sun and their sub-stellar counterparts known as brown dwarfs form in large and complex cosmic clouds of molecular gas and dust. During their formation, the molecular cloud undergoes a hierarchical collapse such that this large cloud shrinks down in size by orders of magnitude, and its mass concentrated into a small central object. This object then evolves to become either a young star or a brown dwarf.

Atomic and molecular outflows in the form of high velocity collimated jets and slower wide angled winds play a critical role during this formation process. Bipolar jets act to remove the excess angular momentum that the central object inherits from its birth environment. Meanwhile, the low velocity winds are thought to remove angular momentum from the circumstellar disc surrounding the central source, hence enabling the accretion of mass into the central object.

The aim of this thesis is to explore unanswered questions in regard to this process. Namely, do these low velocity winds, which are thought to be traced by optical forbidden emission lines actually remove significant angular momentum from the disc, are these winds variable in time? And, what is the role of outflows at sub-stellar masses, particularly at the earliest stages of the formation process? The work presented in this thesis addresses these questions by utilising the forefront observational techniques of optical spectro-astrometry and radio interferometry in order to extract information about the morphology and kinematics of these flows.

This thesis finds that despite variable kinematics, the spatial properties of these low velocity winds in low-mass stars remains stable. Assuming these winds magnetically driven, they don't appear to remove a significant portion of excess angular momentum

from the star-disc system, however, there are substantial uncertainties on these values due to the unconstrained temperature of the outflowing material. Lastly, in the sub-stellar regime, outflows are observed via molecular gas alongside accretion, pointing to a "star-like" formation pathway for the young brown dwarf presented here.

Publications

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List of Abbreviations

1D	one-dimensional
2D	two-dimensional
A_v	Visual extinction
AAVSO	American Association of Variable Star Observers
ACA	Atacama Compact Array
ALMA	Atacama Large Millimetre/sub-millimetre Array
AR	Angular Resolution
ARCs	ALMA Regional Centres
AU	Astronomical Unit
BD	Brown Dwarf
BDs	Brown Dwarfs
BIMA	Berkley-Illinois-Maryland association Array
CALYPSO	Continuum And Lines in Young ProtoStellar Objects
CASA	Common Astronomy Software Applications
CCDs	Charged Coupled Detectors
CRIRES	CRyogenic high-resolution InfraRed Echelle Spectrograph
CMs	Calibrated Measurement sets
CO	Carbon monoxide
CPD	Chromatic Position Difference
CTTS	Classical T Tauri Star
CTTSs	Classical T Tauri Stars
Dec	Declination
DSI	Differential Speckle Interferometry
EHT	Event Horizon Telescope
ELR	Emission Line Region
ELT	Extremely Large Telescope
ESO	European Southern Observatory

ESO's	European Southern Observatory's
ERIS	Enhanced Resolution Imaging Spectrograph
EW	Equivalent Width
FEL	Forbidden Emission Line
FELs	Forbidden Emission Lines
FFT	Fast Fourier Transform
FITS	Flexible Image Transport System
FoV	Field of View
FT	Fourier Transform
FUV	Far Ultra Violet
FWHM	Full Width at Half Maximum
GHz	$\times 10^9$ Hertz
GMC	Giant Molecular Cloud
H₂	Molecular Hydrogen
HDS	High Dispersion Spectrograph
HPBW	Half Power Beam Width
HVC	High Velocity Component
HVC1	High Velocity Component 1
HVC2	High Velocity Component 2
IAU	International Astronomical Union
iFFT	inverse Fast Fourier Transform
iFT	inverse Fourier Transform
IMF	Initial Mass Function
IRAC	InfraRed Array Camera
IRAF	Image Reduction and Analysis Facility
IRAM04191	IRAM 04191+1522
ISM	Interstellar Medium
JAO	Joint ALMA Observatory
JWST	James Webb Space Telescope
LSR	Local Standard of Rest

LVC	Low Velocity Component
LVC-BC	Low Velocity Component - Broad Component
LVC-H	Low Velocity Component - High
LVC-L	Low Velocity Component - Low
LVC-M	Low Velocity Component - Medium
LVC-NC	Low Velocity Component - Narrow Component
MC	Molecular Cloud
MHD	Magnetohydrodynamic
MHz	$\times 10^6$ Hertz
MIRI	Mid InfraRed Instrument
mm	millimetre
MRI	Magnetorotational Instability
MRS	Maximum Recoverable Scale
MuC	Multiple Component
MUSE	Multi Unit Spectroscopic Explorer
MVC	Medium Velocity Component
ngVLA	next generation Very Large Array
NIRSpec	Near InfraRed Spectrograph
nm	$\times 10^{-9}$ meters
OCA	Organization, Classification and Association
OH	Hydroxyl
PA	Position Angle
PAs	Position Angles
PAHs	Polycyclic Aromatic Hydrocarbons
PdBI	Plateau de Bure Interferometer
PLs	Photonic Lanterns
PRT	Photospheric Removal Tool
PSF	Point Spread Function
PV	Position-Velocity
RA	Right Ascension

rms	root mean square
S/N	Signal-to-Noise
SA	Spectro-Astrometry
SC	Single Component
SED	Spectral Energy Distribution
SMC	Small Molecular Cloud
sub-mm	sub-millimetre
THz	$\times 10^{12}$ Hertz
UT2	Unit Telescope 2
UV	Ultra Violet
UVES	Ultraviolet & Visual Echelle Spectrograph
VeLLO	Very Low Luminosity Object
VLT	Very Large Telescope
WTTS	Weak line T Tauri Star
YSO	Young Stellar Object
YSOs	Young Stellar Objects

For those who and inspired me dream big and chase the stars.

Thesis Introduction

From ancient times to the modern age, humanity has strived to understand its origins and place within the Universe. The study of stars has been central to this for millennia and continues to this day. Investigating star formation offers us a window into our very beginnings and the processes by which we have come into being. Over the past several decades the astronomical community has ascertained that the formation of low mass ($\leq 2 M_{\odot}$) stars is a highly dynamic process that takes place over millions of years and has a significant impact on the surrounding interstellar environment. A key part in this process is the launching of outflows from the young forming stars. Since their discovery, astronomers have poured much energy into understanding how outflows are formed, launched, how they affect the young stars themselves and any possible planet formation.

Through these efforts, it is now clear that outflows are ubiquitous structures during the star formation process that seem to always be accompanied by a circumstellar disc of material surrounding the young star. Outflows play a fundamental role in both the star formation and planet formation processes by enabling the accretion of matter onto young stars and contributing to the dispersal and evolution of the circumstellar disc over time. Hence, they have a critical influence on the physical and chemical conditions in the regions where planets may subsequently form. However, despite these advances, there remains several unanswered questions.

This thesis aims to investigate two such remaining questions in regard to outflows utilising the forefront observational techniques of spectro-astrometry and interferometry. Firstly, is the role of outflows in angular momentum transport in the star-disc system. High velocity outflows, known as jets are known to remove a portion of the excess angular momentum that the star inherits from its birth environment during formation,

however, the role of low velocity outflows, known as winds, remains less clear. To address this, this thesis presents spectral and spatial analysis of optical spectra of a small survey of young stars and multi epoch analysis of optical spectra collected primarily with the European Southern Observatory's Very Large Telescope using the technique of spectro-astrometry. These works aim to shed light on the processes which govern the removal of excess angular momentum in these young stellar systems and how these processes vary in time.

The second goal of this thesis is to investigate the role of outflows driven by objects at other masses and to explore the role played by so-called accretion streamers. Brown dwarfs are the intermediary mass objects between low mass stars and giant planets and their primary formation mechanism remains in debate. As such, understanding the role of outflows in the creation process of BDs may unearth key insights into their assembly and investigate possible links between the formation processes of low mass stars and brown dwarfs. This thesis presents interferometric observations at millimetre and sub-millimetre wavelengths taken with the Atacama Large millimetre/sub-millimetre array and provides analysis on the outflows driven by one such unique system.

The thesis is structured as follows:

- *Chapter 1* provides background information on the current understanding of the formation processes of both young low mass and brown dwarfs with a focus on accretion and outflow processes and variability.
- *Chapters 2 and 3* provide an overview of the main techniques used for the collection, reduction and analysis of the data presented in this thesis. Chapter 2 covers the technique of spectro-astrometry, at optical wavelengths. Chapter 3 discusses interferometry at millimetre/sub-millimetre wavelengths.
- *Chapter 4* presents a study on outflow efficiency/angular momentum extraction via the spectro-astrometric analysis of a sample of young low mass stars.

- *Chapter 5* focuses on the spectro-astrometric analysis of a specific star in the collected sample, DG Tau over a period of ~ 20 years to examine what variability may be able to tell us about the origin of the optical forbidden emission line low velocity component.
- *Chapter 6*, presents the interferometric analysis of the youngest, lowest mass object presented in this thesis, IRAM 04191+1522 to study the role out outflows at masses below the stellar mass regime.
- *Chapter 7* finishes the thesis, providing overall conclusions and a discussion on the avenues of future research.

Chapter 1

Background

Jets and winds, collectively referred to as outflows, are an ever present feature of low-mass young stars and are thought to play a major role in the star formation process. Observations of outflows offer critical insights into not just the mass assembly processes in young stars but, also how outflows shape the environment in which young planets may form. In addition, outflows eject fresh material into the surrounding interstellar medium, changing the physical and chemical conditions in regions where new generations of stars are born. In conjunction with this work in the stellar mass regime, observations have also shown that outflows are present during the formation of brown dwarfs, the sub-stellar mass "cousins" to young stars. These revelations now raise interesting questions about the role of outflows during the formation process of brown dwarfs and if that role is comparable to the role of outflows in the formation of young stars.

Advances in forefront observational techniques, such as high resolution spectroscopy, spectro-astrometry and interferometry now offer the ability to study the kinematics and morphology of outflows launched by both young low-mass stars and brown dwarfs in unprecedented detail in close proximity to the outflow launching region. These kinds of observations provide critical information required to test current theoretical models of outflows. This chapter provides the background information concerning the current understanding of low-mass star and brown dwarf formation.

This chapter begins with a discussion on the process of low-mass star formation. Following this is an explanation of the accretion processes in these stars, a short review of the main outflow launching models and an introduction to the angular momentum

transport problem. The remaining sections of this chapter will focus on explaining the main branches of brown dwarf formation and also the observed variability present in both the mass accretion and mass outflow processes.

1.1 Low-mass star formation.

Broadly speaking, the formation process of young low-mass stars can be divided into six progressive stages, from a molecular cloud (MC) core to a fully formed stellar system as outlined in Fig. 1.1. Each stage in the star formation process is dominated by the collapse of sparse, large scale structures comprised of gas and dust into progressively smaller scale structures where the mass becomes more concentrated. Below, the stages of low-mass star formation are outlined in more detail.

1.1.1 The birth of a low-mass protostar

It is generally accepted that low-mass ($\leq 2 M_{\odot}$) young stellar objects (YSOs) form within MCs, composed of cold and dense of molecular gas and dust within the interstellar medium (ISM; Lada, 1987). MCs are gravitationally bound over dense regions of the ISM and can be formed by various processes. For example, MCs can form from shock wave induced turbulence from supernovae explosions (Zucker et al., 2022), compression via galaxy spiral density waves (Dobbs, Bonnell, and Pringle, 2006; Stark and Lee, 2006) among other processes (see Dobbs et al., 2014, and references therein). MCs can be divided into two types, Small Molecular Clouds (SMCs) and Giant Molecular Clouds (GMCs). These types of MCs are defined based on the mass within the cloud and how far it spans. SMCs span up to a few parsecs in size and range in mass up to $\sim 10^2 M_{\odot}$. GMCs on the other hand can be up to hundreds of parsecs in size with masses ranging from $10^4 - 10^6 M_{\odot}$ (Hartmann, 2009; Ray et al., 2007). Most of our understanding of low-mass star formation comes from studies of MCs within the Milky Way and as such, the remainder of this discussion focuses on star formation within these objects.

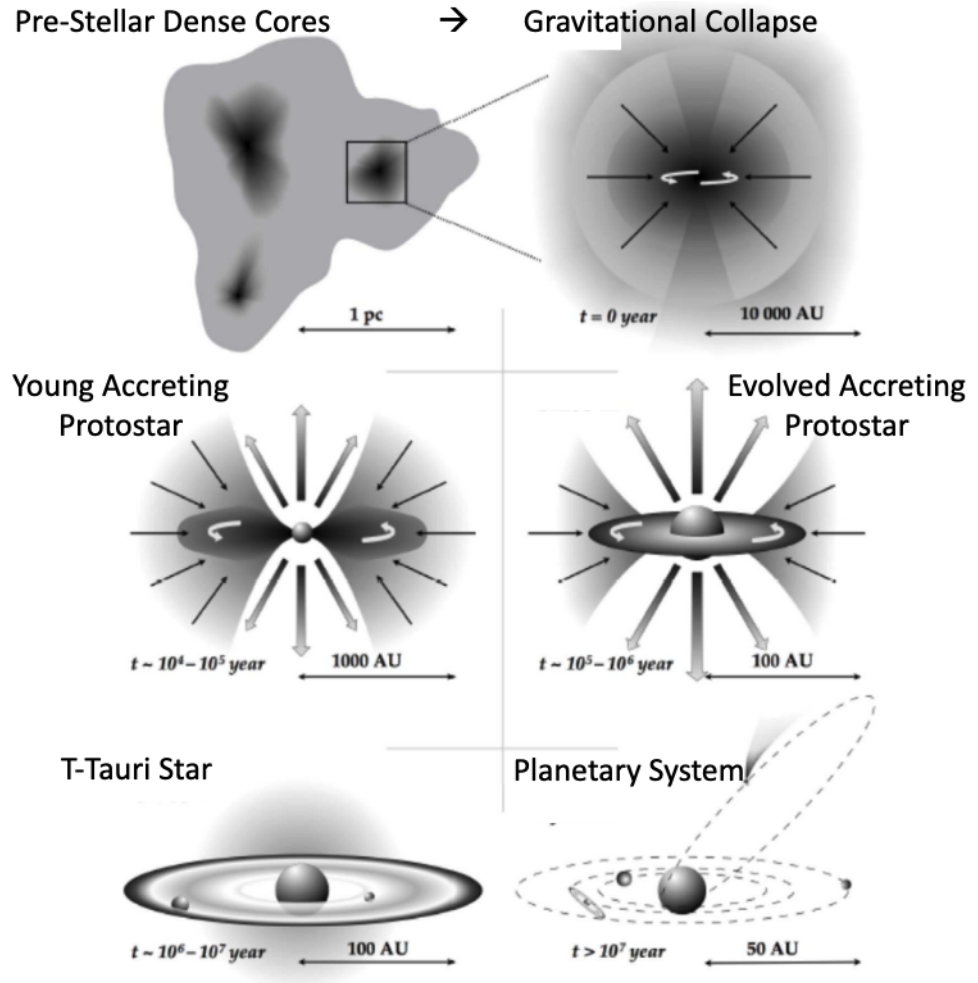


FIGURE 1.1: Cartoon of the classical picture of low-mass star formation after the formation of dense cloud cores. (Colzi, 2020).

Most of the mass contained within a MC is in the form of molecular hydrogen (H_2) with a mean density of $n_{\text{H}_2} = 100 \text{ cm}^{-3}$ (Dame, Hartmann, and Thaddeus, 2001). It is difficult to observe H_2 as it rarely emits at the ~ 10 to 100 K temperatures within these clouds because the rotational transitions of H_2 require high temperatures (several 100s of Kelvin) to be excited (Valdivia, Valeska et al., 2016). However, MCs do contain some atomic hydrogen that is generated by the photodissociation of H_2 by Far UltraViolet (FUV) photons from background stellar populations, and trace amounts of dust and carbon. Some H_2 may be shielded from photodissociation via molecular self shielding and by the absorption and remission of some FUV photons by the dust. These processes allow for the eventual formation of carbon monoxide (CO) molecules. Specifically, CO molecules form via reaction chains of ion–neutral species or neutral–neutral

species (see Sternberg and Dalgarno, 1995, for a more detailed description of CO formation chemistry). In essence, cosmic rays penetrate the outer regions of the MC and ionise some of the atomic and neutral hydrogen and carbon to form their ionised species H^+ and C^+ . These ions react with hydroxyl (OH) groups¹ to form CO. The main chemical formation route of CO in a MC is shown in Equation 1.1 (van Dishoeck and Black, 1988).



CO and its less abundant isotopologues (e.g. ^{13}CO , $C^{18}\text{O}$) have been used for decades as tracers of dense gas within MCs (Dame, Hartmann, and Thaddeus, 2001; Dame et al., 1987; Heyer and Dame, 2015) as it can act as a proxy to H_2 (Bolatto, Wolfire, and Leroy, 2013). From observations of CO, it has become clear that the density and velocity structure of these clouds is not uniform but, rather is composed of a hierarchical network of filamentary structures (e.g. Hacar et al., 2023; Pineda et al., 2023, and see Fig. 1.2).

These structures are produced by various mechanisms including gravity, propagating shock waves, heating, cooling, magnetic fields and chemistry within the MC and the subsequent gravitational fragmentation (Chevance et al., 2023, and references therein) and/or turbulence driven fragmentation (Ishihara et al., 2025) of these filaments form cloud cores. Cloud cores are localised, high density regions (i.e. the molecular hydrogen number density, $n_{H_2} \geq 10^4 \text{ cm}^{-3}$; Hacar et al., 2023) within MCs from which, young low-mass stars can be born. The exact mechanism through which MCs collapse to form filaments and cloud cores remains in debate but, the literature proposes two scenarios, either a hierarchical collapse of small local regions within the cloud (Vázquez-Semadeni et al., 2019) or a global contraction and collapse of the cloud up to scales of the cloud itself (Heitsch, Hartmann, and Burkert, 2008; Heitsch et al., 2008).

¹Oxygen is released into the gas phase when cosmic rays sublimate the water ice that is frozen in the dust grains at the low temperatures of the MC. This released gas phase oxygen reacts with atomic hydrogen to form OH groups.

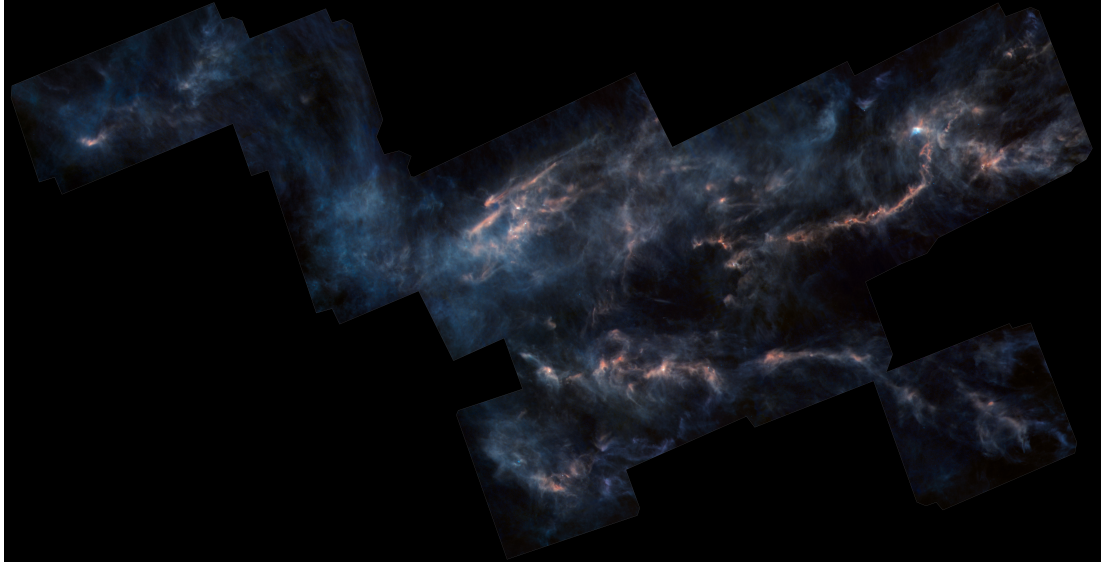


FIGURE 1.2: Herschel space telescope four colour mosaic of the Taurus molecular cloud located ≈ 140 pc away. The image combines observations at 160 microns (blue), 250 microns (green), 350 microns (divided between green and red) and 500 microns (red). This image shows that this molecular cloud consists of a network of filamentary structures of gas and dust. The image spans 13.8° by 7.3° . North is up and east is to the left. (Credit: ESA/Herschel/NASA/JPL-Caltech; acknowledgement: R. Hurt (JPL-Caltech)).

Observational evidence (e.g. Chevance et al., 2020; Kruijssen et al., 2019) strongly suggests that the efficiency at which the MC material is converted into stars is of the order of a few percent. This raises questions about both collapse scenarios. At present, there is no consensus on a dominant mechanism for cloud collapse. Irrespective of formation mechanism, once a cloud core is formed, it remains stable provided that the internal kinetic energy (E_k) of the particles in the core generate a thermal pressure strong enough to balance the self-gravitational potential (U) as outlined by the virial theorem (Equation 1.2).

$$2E_k + U = 0 \quad (1.2)$$

The cloud core will continue to accrete material from its parental surroundings until it reaches a critical mass, referred to as the Jeans critical mass (M_J ; Jeans, 1902), such that thermal pressure can no longer balance the gravitational forces. In the simplest case of an isothermal, initially uniform cloud core with density ρ_0 , pressure P and

sound speed c_s , the Jeans critical mass is defined as shown in Equation 1.3. Where G is the universal gravitational constant. A full treatment of this derivation is provided by several sources (e.g. Hartmann, 2009; Shu, 1977).

$$M_J = \left(\frac{\pi c_s^2}{G} \right)^{3/2} \rho_o^{-1/2} \quad (1.3)$$

Assuming a core consisting of a gas with a mean molecular weight of 2.36^2 gives Equation 1.4. Where T_{10} is the temperature of the gas in units of tens of Kelvin and n_{H_2} is the H_2 density in cm^{-3} . Typical cloud core masses are in the range of a fraction of a solar mass to $\sim$ a few hundred $M_\odot$ with Jeans masses in the nearby star forming regions ranging from ~ 0.5 to a few $M_\odot$'s (see Motte et al., 2022, and references therein).

$$M_J \approx 540 M_\odot T_{10}^{3/2} n_{H_2}^{-1/2} \quad (1.4)$$

When the Jeans mass is reached, the cloud core begins to collapse. This initial collapse leads to the formation of the so-called first hydrostatic core (Larson, 1969). This short-lived core is mostly still molecular with optically thick dust continuum emission. This first core continues to grow by accreting mass from the continued layers of infalling material (Boss and Yorke, 1995; Omukai, 2007). As the first hydrostatic core grows, the effect of radiative cooling is minimised as the central region is optically thick, hence, when temperatures inside the core reach approximately 2000 K; H_2 molecules begin to dissociate into atomic hydrogen via collisional excitation. This conversion of H_2 into atomic hydrogen removes thermal energy from the system and gravitational forces again begin to dominate. This leads to a further collapse of the first hydrostatic core. As the first hydrostatic core undergoes this second collapse, about half of the gravitational energy is converted into kinetic energy according to the virial theorem (see Equation 1.2) and the rest is converted into thermal energy via collisions between particles (i.e. frictional heating). This friction heats the central region until temperatures of $\sim \times 10^6$ K are reached, when the burning of deuterium begins. This leads to the ignition of the central region into the true protostar, with enough thermal pressure to

²This number is derived assuming a mixture of gas consisting of $\sim 82\%$ H_2 , $\sim 18\%$ helium and trace amounts of heavier elements (Padovani, Galli, and Glassgold, 2009).

resist further contraction via the self gravitational potential. The protostar remains surrounded by the bulk of the leftover parental core, which is referred to as the envelope.

Due to the conservation of angular momentum, rotation of the initial parental core leads to the formation of a flat, centrifugally supported circumstellar accretion disc (Inutsuka, 2012; Terebey, Shu, and Cassen, 1984). From this accretion disc, the newly formed protostar accretes matter to grow and evolve (Hartmann, Herczeg, and Calvet, 2016; Shu, Adams, and Lizano, 1987). The evolutionary stages of low-mass stars are discussed further in Section 1.1.2. In addition, magnetic fields and rotation lead to the production of bipolar jets and winds (Pudritz et al., 2007; Ray et al., 2007). There are two competing models explaining the production of jets and winds: the X-wind and disc wind models, which are discussed further in Section 1.2. These jets and winds in turn drive molecular outflows (Arce et al., 2007; Santamaría-Miranda et al., 2020; Wu et al., 2004).

1.1.2 Protostellar Evolution

After the young protostar is formed, in the classical picture of low-mass star formation, the star-disc system follows a chain of evolving stages on its path to becoming a fully formed stellar system. These steps are outlined in Fig. 1.3. Class I, II and III YSOs are defined based on the spectral index (α) measured from near to mid infrared part ($\sim 2 - 20 \mu\text{m}$) of the Spectral Energy Distribution (SED) of the star (Lada, 1987) and/or on the bolometric temperature (Chen et al., 1995; Evans et al., 2009). The spectral index is essentially the slope of the SED in this wavelength range as shown in Equation 1.5. Where S_λ is the flux density at a specific wavelength λ (Dunham et al., 2014).

$$\alpha = \frac{d \log(\lambda S_\lambda)}{d \log(\lambda)} \quad (1.5)$$

Evolution of young stellar objects

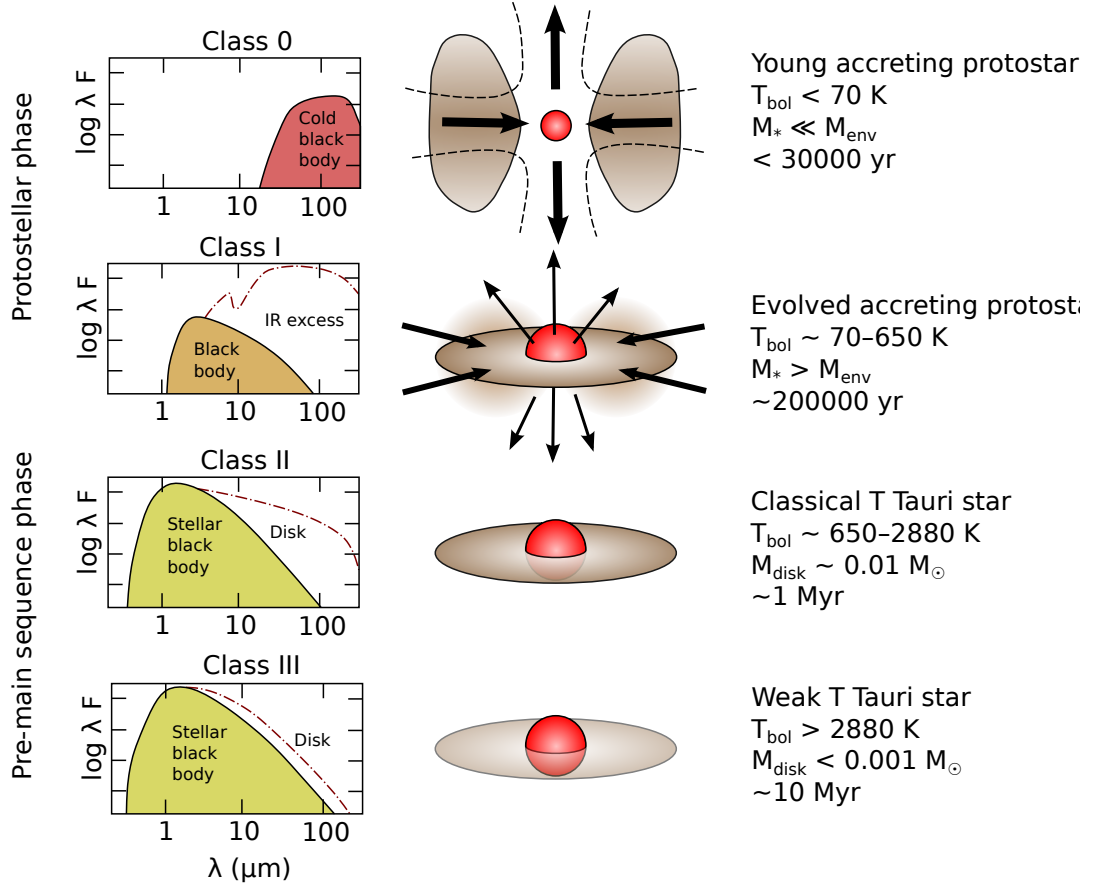


FIGURE 1.3: Stages of protostellar evolution with the corresponding SED. The central object is represented as the blackbody contribution to the SED that is covered in an infrared excess. The peak of the blackbody shifts and the amount of infrared excess changes as the central object and the surrounding environment evolves. Figure adapted from Pols, 2011.

Class 0

Class 0 protostars are only observed at millimetre or radio wavelengths as they are heavily obscured by their birth environments, with visual extinctions (A_V) of ≥ 50 (Laos et al., 2021). They represent the earliest stage of protostellar evolution where over 50% of the systems mass is still contained within the natal envelope surrounding the young protostellar core (Stock et al., 2020). These objects have a core mass of $\sim 1 M_{\odot}$. Class 0 objects as a subtype of low-mass YSOs were added later than the other stages to the evolutionary track shown in Fig. 1.3, as they are extremely submerged in

their optically opaque environments, making them undetectable in the near infrared (Greene et al., 1994). Their SEDs are dominated by the thermal emission of the gas and dust surrounding the core as the core itself has a very low intrinsic luminosity. Class 0 objects are typically found via other means than direct observation. For example, as these objects are vigorously accreting mass, they drive large outflows (see Fig. 1.4). Through observations of large outflows in warm ($\sim$ few 100 K), shock tracing molecular emission lines, such as SiO, Class 0 YSOs can be detected (Andre, Ward-Thompson, and Barsony, 1993; Podio et al., 2015, 2016).



FIGURE 1.4: A JWST NIRCам three colour composite image (taken from Ray et al., 2023) of the bipolar outflow emitted from the young ($\text{few } \times 10^4$ years) star, HH211. The three filters are centred at $3.35\mu\text{m}$ (blue), $4.60\mu\text{m}$ (green) and $4.70\mu\text{m}$ (red). The outflow is redshifted to the northwest and blueshifted to the southeast. These filters contain several of H_2 and CO lines known to trace outflowing gas. The star itself is obscured by the disc, which is the dark central lane in between the jet lobes in this image.

Class I

As the protostar continues to accrete material from the envelope, it heats up the surrounding dust and gas hence, it becomes observable in the near infrared as the spectral index shifts to $\alpha \geq 0.3$ (Greene et al., 1994). Class I objects remain obscured by their birth environment. The SED can be used to estimate how much of the natal envelope has been cleared away with the transition from Class 0 to Class I, being defined by either a bolometric temperature ≥ 70 K or by the amount of intrinsic bolometric luminosity that is observed in the sub millimetre (Evans et al., 2009). These objects drive strong, very collimated and extended jets which are thought to aid in the removal of the excess angular momentum the protostar inherits from its parental environment (Caratti o Garatti et al., 2024; Podio et al., 2015; Ray et al., 2023).

Class II

The Class II stage represents the transition of the central object (which is now $\leq 2 M_{\odot}$) from its last phase as a Class I protostar to its first phase as a pre main sequence star (André et al., 2004). In the Classical T Tauri star (CTTS) stage, most of the surrounding material from the natal envelope has been ousted from the star-disc system through a combination of stellar mass accretion, irradiation and stellar outflows (i.e. jets and winds). The outflows driven by Class II YSOs are significantly less powerful than in previous stages. Given this, the mass accretion rate drops significantly compared to previous stages as solely the circumstellar accretion disc is feeding the star. The remaining envelope becomes optically thin and the pre main sequence star at the centre of this system becomes visible at optical wavelengths. The spectral index of these stars is between $-1.6 \leq \alpha < -0.3$ (Greene et al., 1994) as the slope of the SED is decreasing. The infrared excess observed in this stage is dominated by the presence of the circumstellar disc. It is in these discs that one could search for evidence of planet formation as this is the last evolutionary stage where the disc has a significant mass (Lada, 1987; Testi et al., 2022).

The spectra of these stars show a forest of permitted and forbidden emission lines (Hartmann, 2009; Whelan, 2014) that trace both strong accretion (e.g. through the Balmer series of hydrogen lines, $H\alpha$, $H\beta$ etc.) and ejection processes (e.g. traced by optical forbidden emission lines). Objects which present a "Flat SED" (spectral index of between $-0.3 < \alpha < 0.3$ in the mid infrared; Greene et al., 1994) have a slope which distinguishes them from the typical SEDs of Class I and II objects. This makes them a possible transitional stage between these two classes of YSO.

Class III

This stage represents the final part of low-mass YSO evolution before the star reaches the main sequence. By this stage, the spectral index of the system has shifted to $\alpha < -1.6$, natal envelope is completely dispersed and accretion and ejection has practically halted. These stars are also referred to as Weak Line T Tauri stars (WTTS) as their spectra are absent of the strong accretion tracing lines present in the spectra of CTTS. For WTTSs the equivalent width of the $H\alpha$ is typically measured to be $< 10 \text{ \AA}$ (Appenzeller and Mundt, 1989). Additionally, the circumstellar discs around these objects have evolved significantly, showing evidence of dispersal and substructures (Hughes et al., 2025; Lovell et al., 2026) which may be caused by a variety of processes, including planet formation (Capelleveen et al., 2025; Keppler et al., 2018).

1.2 Accretion and Ejection

1.2.1 Angular momentum transport

The young stellar systems described above inherit the majority of their angular momentum from the massive parental environments during collapse. In order for these young stars to grow and evolve, this excess of inherited angular momentum must be removed, so that the star does not spin extremely rapidly to the point of breaking itself apart. This is alleviated through the launching of bipolar outflows (jets and winds)

which aid in the removal of this excess angular momentum (see Section 1.2.4) and regulates the stellar rotation (Frank et al., 2014; Ray and Ferreira, 2021).

Additionally, the stellar magnetic field acts as a "break" to the disc itself (Liffman, 2026). Angular momentum must also be transported outward from the circumstellar disc, to allow for the accretion of material from the inner disc onto the star (see Section 1.2.2.) This was initially thought to be handled by the so-called magneto rotational instability (MRI). The MRI is a type of disc instability initially proposed by Balbus and Hawley, 1991 to explain the radial transport of excess angular momentum in a viscous accretion disc and the corresponding observed stellar mass accretion rates in YSOs. Figure 1.5 presents a schematic diagram showing how the MRI is induced, and this process is now briefly described. A deeper treatment of the MRI and viscously driven angular momentum transport is presented in the review article of Balbus, 2003. In short, the instability arises via the interaction of differentially rotating material in the disc with magnetic tension via magnetic field lines threading the disc. Given two "packets" of disc material m_o and m_i (for outer mass and inner mass respectively) separated by an initial radial displacement, where m_i is closer to the central mass (M_c ; i.e. the star) and therefore moving at a larger angular velocity; the poloidal magnetic field applies a tension to m_o and m_i such that the field acts like a spring applying a restoring force to m_o and m_i . However, as the disc is rotating, the magnetic tension applied by the poloidal field results in the transport of angular momentum outward from m_i to m_o . Therefore, the angular velocity of "packet" m_o increases and the angular velocity of "packet" m_i decreases. Hence, "packet" m_i moves towards the central mass and "packet" m_o moves out to another part of the disc. This motion of differentially rotating "packets" of mass stirs material in the disc and from this, the MRI was thought to drive the disc turbulence necessary to transport excess angular momentum away from star-disc system and enable stellar mass accretion.

However, more recent simulations (Bai, 2016; Bai and Stone, 2013) have shown that the effect of the MRI is suppressed in regions of the disc beyond a few 10s of AU

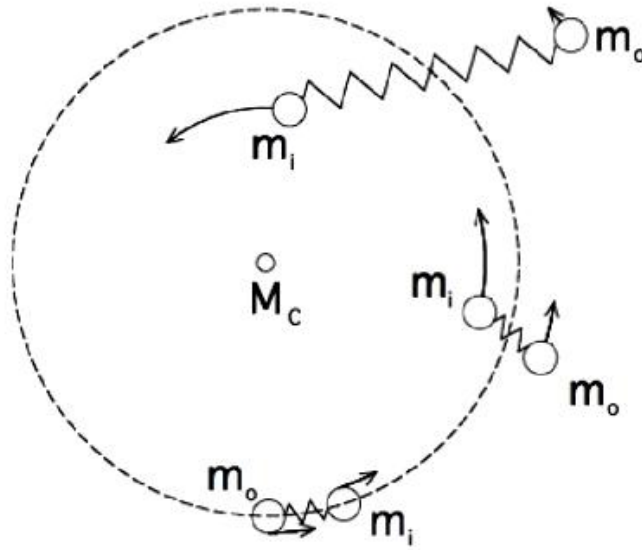


FIGURE 1.5: Schematic diagram showing how the MRI is generated.
Taken from Balbus, 2009.

due to insufficient disc ionisation and non-ideal magnetohydrodynamic (MHD) effects (see Pascucci et al., 2023, and references therein). Additionally, Atacama Large millimetre/sub-millimetre Array (ALMA) observations have shown that the turbulence generated by the MRI in regions of the outer disc is likely quite weak (Flaherty et al., 2015). The lack of a reasonable mechanism to explain the transport of excess angular momentum in circumstellar discs at large radii has led to the so-called angular momentum transport problem in star formation.

1.2.2 Magnetospheric accretion

The following sections describe the accretion mechanisms for YSOs, outline the main outflow launching models (the X wind and disc wind models) and explain how a renewed interest in optical forbidden emission lines (FELs) may shed light on the angular momentum transport problem.

Magnetospheric accretion is generally accepted to be the process which governs the transfer of mass from the circumstellar disc to the central YSO in the low-mass regime ($< 2 M_{\odot}$). This process is depicted in Figure 1.6 (taken from Hartmann, Herczeg, and Calvet, 2016). The transport of mass from the circumstellar disc to the star is governed

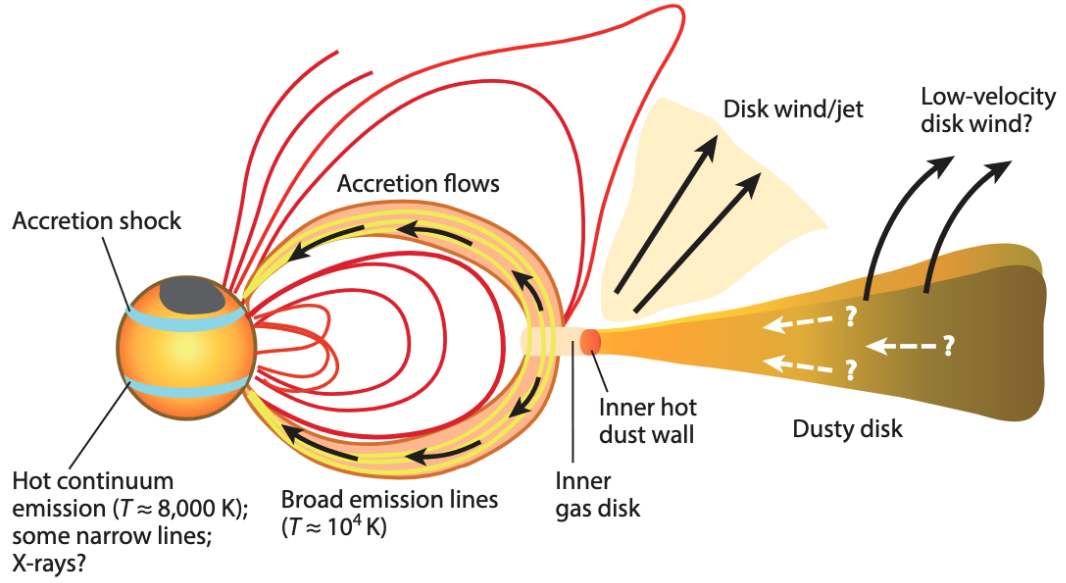


FIGURE 1.6: Schematic diagram of the magnetospheric accretion process taken from Hartmann, Herczeg, and Calvet, 2016.

by the interaction between the stellar magnetic field and the inner region of the disc (Bouvier and Appenzeller, 2007). Material from the circumstellar disc is moved inward toward the inner edge of the disc, provided it loses angular momentum through some mechanism. For example, these may include gravitational instabilities (Kratte and Lodato, 2016), the Hall effect (see Wardle and Salmeron, 2012, and references therein) or disc winds (Gressel et al., 2015; Pudritz and Norman, 1983). The inner disc edge is created due to the truncation of the circumstellar disc by the dipole stellar magnetic field that threads through the disc. This truncation radius is typically $\sim$ a few stellar radii from the stellar surface (Pudritz, Ralph E. and Matt, Sean, 2014).

Magnetospheric accretion happens at this inner truncation radius. Stellar photons heat and ionise material that is to be accreted. The temperature of disc material at the truncation radius is $> 10^3$ K, so the dust in this region sublimates. These temperatures are high enough to allow for strong coupling between the sublimated gas and the stellar magnetic field (Koenigl, 1991). Magnetic pressure generated from the stellar magnetic field acts on the inner disc edge. Gas ram pressure from the surrounding disc also acts on the material at the inner disc edge. When this magnetic pressure overcomes

the gas ram pressure, the ionised material at the inner disc edge is directed towards the star along the stellar magnetic field lines. This forms accretion columns that move material from the inner disc edge to the stellar surface. At the disc surface the accreted gas is temporally shock heated to temperatures up to 10^6 K, which causes hotspots to form on the stellar photosphere where these accretion columns hit the stellar surface.

The magnetospheric accretion model is evidenced by a plethora of observations. Calvet and Gullbring, 1998 argue that the observed "filling in" of photospheric absorption features in the spectra of CTTSs is due to an excess in optical/ultraviolet (UV) emission caused by the accretion shock when the accreted material hits the stellar surface. Additionally, Edwards et al., 1994 observe optical and near infrared redshifted absorption features with velocities comparable to the free fall velocity of infalling material (~ 100 to a few 100 km s^{-1}). Lastly, the observed photometric variability of CTTSs may also be partly attributed to accretion hotspots modulated by stellar rotation (Alencar et al., 2012; Bouvier et al., 1999, 2007; Gullbring et al., 1996).

1.2.3 Accretion streamers

Alongside magnetospheric accretion, recent observations of so called accretion streamers paint a picture of a process whereby "fresh" material may be fed to the circumstellar disc. The detection of these structures adds complexity to this classical model of isolated star formation, as these structures have been observed to originate outside the envelope (Pineda et al., 2020). Accretion streamers are elongated, asymmetric structures of gas observed primarily at sub-millimetre and millimetre wavelengths (Fig. 1.7). They act as a bridge between the circumstellar disc and its large scale natal environment.

In recent years they have been observed around various different protostars and pre-main sequence stars (e.g. Artur de la Villarmois et al., 2026; Hales et al., 2024; Hsieh et al., 2023; Liu et al., 2025; Pineda et al., 2020; Yen et al., 2019) in several molecular emission lines. They often exhibit varying lengths and smooth velocity gradients,

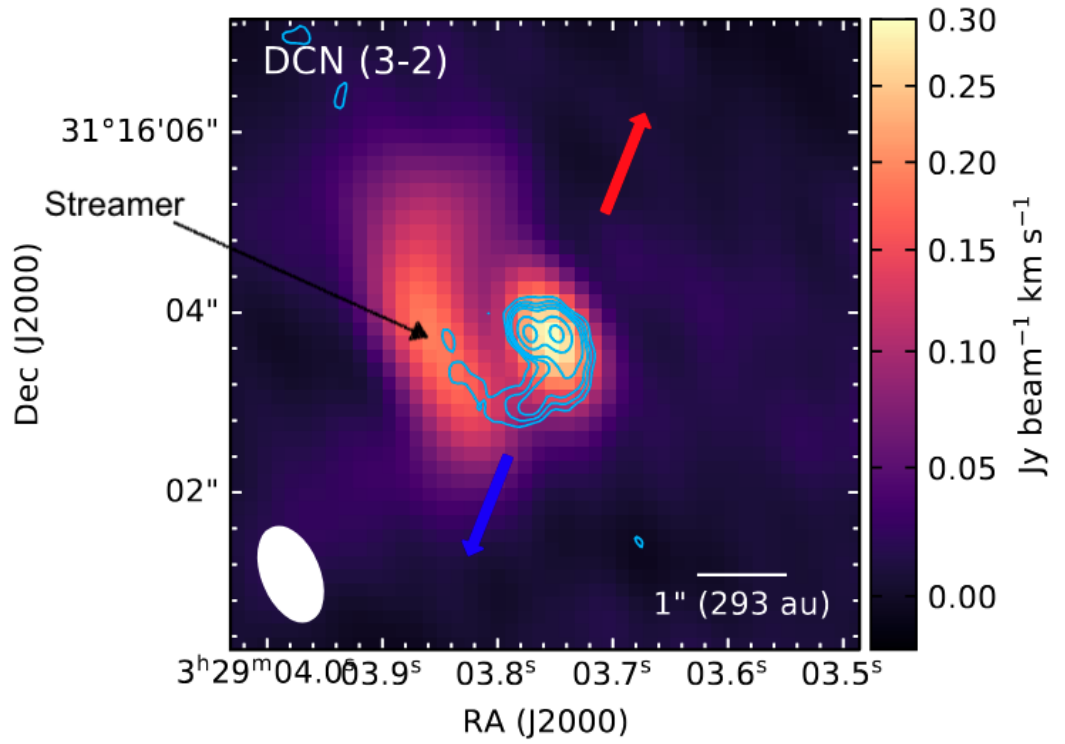


FIGURE 1.7: Example of an accretion streamer feeding Class 0/I binary system SVS13A, taken from Hsieh et al., 2023. Shown is the integrated intensity map of the molecule DCN from 8.26 to 8.8 km s⁻¹, overlaid on this is the 1.3 mm continuum emission shown in blue contours. The red/blue arrows indicate the directions of the outflow traced in CO emission (Lefèvre et al., 2017).

seemingly indicative of an infalling flow towards the circumstellar disc and/or central star. The impact of accretion streamers on the star formation process is still relatively unexplored however, this section aims to outline some of their potential impacts on the star-disc system and the potential implications on the star formation process.

Streamers provide chemically unprocessed gas to the circumstellar disc (Tanious et al., 2025). By feeding the disc, streamers could impact the physical and chemical properties of discs during star and planet formation (Taniguchi et al., 2024). Additionally, providing the disc with new gas may also affect the type and amount of material available to the YSO for accretion. The impact site of streamers is also of interest. Streamers can be traced via shock tracing molecular emission lines (e.g. sulphur bearing species, Artur de la Villarmois et al., 2026). These shocks heat the impact region of the disc. Streamers may also induce disc instabilities (Flores et al., 2023) which may subsequently contribute to the creation of disc substructures. Recent simulations (Hühn and Dullemond, 2025) and growing observational evidence (Garufi et al., 2022; Hanawa et al., 2024) also suggest that Class II discs may experience significant interaction with their surrounding environment via accretion streamers.

Many questions regarding streamers remain, estimates have been made of their total mass and lifetimes (Taniguchi et al., 2024). However, the typical mass and lifetimes are poorly constrained due to a limited number of observations. In addition, there are questions regarding the physical and chemical structure of streamers themselves and uncertainty about how much mass may be delivered to the star-disc system over any given period of time (Nakamura et al., 2024).

1.2.4 Magnetohydrodynamic wind models

There are two competing models that describe how outflows are magnetocentrifugally driven by young stars and their circumstellar discs. These two models are the "X-wind" and "disc-wind" (or D-wind) models, initially proposed by Shu et al. (1994) and Shu et al. (1988) and Pudritz and Norman (1983, 1986) respectively. The D-wind

model in particular is heavily influenced by the work of Blandford and Payne (1982). Blandford and Payne (1982) show that an outflow can be generated from the disc if the poloidal component of the magnetic field makes an angle with the disc surface below some critical angle. The following sections provide a brief description of these models. An in-depth review of these models can be found either in the aforementioned papers or in the review article by Ray and Ferreira, 2021.

The X-wind model

In the X-wind model, the outflows arise from the interaction between the stellar magnetic field and the inner circumstellar disc at a narrow annulus or "X-point" (see Fig. 1.8).

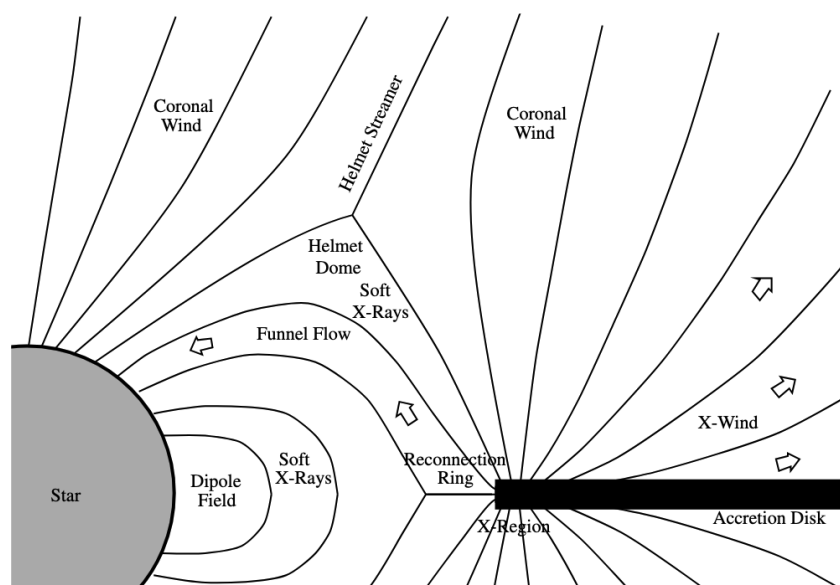


FIGURE 1.8: A diagram of the X-wind model taken from Shu et al., 2000. In this model, the stellar magnetic field truncates the disc and ionised material is launched along open magnetic field lines to generate the outflow(s).

The model outlines that the stellar magnetic field truncates the surrounding disc at an inner co-rotation radius where the Keplerian rotation rate of the disc equals the stellar rotation rate. At this radius, the stellar magnetic field lines that are anchored to the disc, begin to "twist" due to the difference in angular velocity between the inner and outer disc. This causes the formation of accretion flows to the star (see Section

1.2.2 and Hartmann, Herczeg, and Calvet, 2016) and causes part of the magnetic field that threads the disc beyond this radius to become diffuse and switch to an "open" configuration (Ray and Ferreira, 2021). Additionally, the close proximity of the young star to the inner disc implies that the material in this inner disc region is ionised by the star and there is strong coupling of the ionised gas to the magnetic field (Koenigl, 1991). This ionised gas is then centrifugally accelerated along the open magnetic field lines to form an outflow, which transports angular momentum away from the star-disc system. The inner part of this flow becomes more collimated to form a high velocity jet as the rest of the ejected flow becomes a slow and wide angled wind. The predicted jet ejection radius is in the order of $\simeq 0.5$ au (Bjerkeli et al., 2016; Pascucci et al., 2023).

The D-wind model

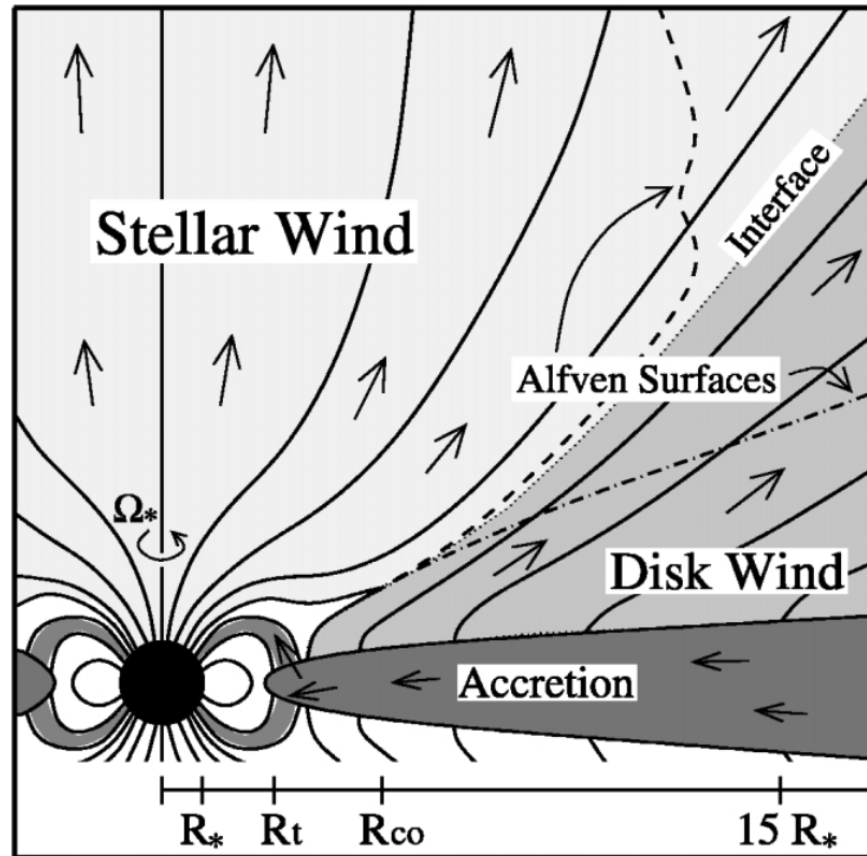


FIGURE 1.9: A diagram showing the D-wind launching model. Like the X-wind model the stellar magnetic field truncates the surrounding disc but conversely, the outflow(s) are launched over a large range of radii. Taken from Matt and Pudritz, 2005.

The most important distinction of the D-wind model compared to the X-wind model is that the jets and winds are launched over a range of radii ($\sim 0.3 - 4$ AU) rather than being launched from a single narrow region of the disc (see Fig. 1.9). This model proposes that the poloidal component of the stellar magnetic field lines is spread throughout the disc at various radii. From this field, the high velocity jets are launched from the innermost radii of the disc in a self collimating manner (Blandford and Payne, 1982), similar to the X-wind model. The slower, wider winds are launched via a spin-down torque exerted on the magnetized rotating disc. They are then launched along field lines anchored at a wide range of disc radii. The D-wind model predicts a continuum of radii and launch velocities, where the velocity decreases with increasing radius (Konigl and Pudritz, 2000; Pudritz et al., 2007) leading to a layered "onion like" outflow structure. Recent observational evidence supports the D-wind model through spatially resolved observations of these radially stratified outflow structures (Birney et al., 2024a; de Valon et al., 2022; Pascucci et al., 2024).

1.2.5 Optical forbidden emission lines

Optical forbidden emission lines are spectral lines that may be produced in the low density environments of space. They are forbidden due to the fact that these transitions break one or more of the standard selection rules of quantum mechanics and as such, have very low transition probabilities. In addition, they have long radiative decay times (~ 100 s of seconds) and hence, require low density environments (e.g. the outflows launched by CTTs) to avoid collisional de-excitation.

From the earliest studies of the optical spectra of CTTs (Edwards et al., 1989; Hartigan, Edwards, and Ghandour, 1995), it was noted that the FEL profiles of many of these objects displayed multiple velocity components. These early studies found two distinct velocity components. A spatially extended High Velocity Component (HVC, $|\nu| \geq 100 \text{ km s}^{-1}$) and a spatially compact Low Velocity component (LVC, $|\nu| \leq 40 \text{ km s}^{-1}$). The spatially extended HVC has long been associated with the collimated jet launched by YSOs (Hirth, Mundt, and Solf, 1997), whereas the origin of the LVC

remains a matter of intensive study (e.g. Banzatti et al., 2019; Fang et al., 2018; McGinnis et al., 2018; Nisini et al., 2024; Simon et al., 2016).

The LVC has been speculated to be tracing a wind (Kwan and Tademaru, 1988), however the exact "flavour" of wind it traces is unknown. Previous studies have shown that this wind may be magnetic in nature, tracing an MHD wind (Birney et al., 2024a; Pascucci et al., 2023) or thermal in nature and therefore tracing a photoevaporative wind driven by the radiation emitted by the young star (Rab et al., 2023). This distinction is critically important as the renewed interest in the study of the LVC in recent years stems from the possibility that the MHD winds, traced by the LVC, are the key to solving the angular momentum transport problem in the current models of low-mass YSO formation (see Section 1.2.1).

If the LVC does in fact trace an MHD wind, this has massive implications for not only stellar mass assembly, but also for disc dispersal and ultimately the planet formation process. For example, the simulations of Bai (2016, 2017), which include non-ideal MHD effects (e.g. ambipolar diffusion), have shown that low velocity winds in the inner circumstellar disc provide the spin-down torques required for the removal of excess angular momentum and hence, enable stellar mass accretion. These models consequently predict that the mass outflow rate for the wind should be comparable to the stellar mass accretion rate. The authors arrive at this prediction as their simulations show that an increasing penetration depth of external (i.e. from the central star) FUV photons into the circumstellar disc sufficiently ionises the disc in deeper layers closer to the disc midplane. This allows for efficient coupling of the gas to the field lines. This effect of external photon heating then implies that the wind becomes more heavily loaded with material (to be ejected) and hence, this rapidly increases the wind mass loss rate to values that are approximately comparable to the observed stellar mass accretion rate in CTTs. This prediction can be tested observationally. Therefore, finding a reliable tracer of these winds is of critical importance to test these simulations and in understanding angular momentum transport in YSOs. The FEL LVC is a

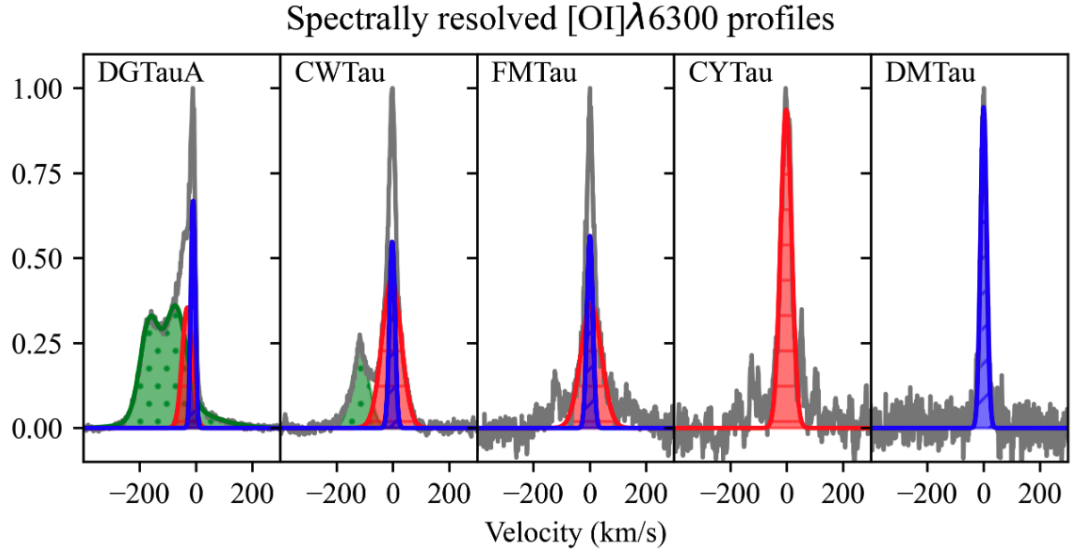


FIGURE 1.10: Example of multi-component optical FEL profiles taken from Pascucci et al., 2023. Green = HVC, Red = LVC-BC and Blue = LVC-NC.

promising potential tracer of these low velocity winds and as such, may be important in the removal of excess angular momentum in these young stellar systems. However, the compact nature of the LVC presents an observational challenge to understanding its origin. The first of this new wave of studies on the LVC used the technique of kinematic fitting via Gaussian decomposition to address the compact nature of the LVC and study it. These works are now described briefly.

Rigliaco et al. (2013) and Simon et al. (2016) expand on these earliest spectroscopic studies by using much higher spectral resolution data to study the optical FELs of many CTTS. They find that the LVC can be further subdivided into two distinct sub-components. They refer to the broad and narrow sub-components as the LVC-BC and LVC-NC respectively. Simon et al. (2016) bases the distinction between these sub-components on the Full Width at Half Maximum (FWHM) measured from the kinematic fitting of each sub-component. The LVC-BC has a $\text{FWHM} > 40 \text{ km s}^{-1}$, whereas the LVC-NC has a measured $\text{FWHM} \leq 40 \text{ km s}^{-1}$ (Fig. 1.10). The rationale for this cut off in FWHM between the LVC-NC and LVC-NC comes quantitatively from the distribution of FWHM values in the sample of $[\text{O I}] \lambda 6300$ line profiles presented in Simon et

al. (2016). In their sample 30 [O I] $\lambda 6300$ line profiles; the FWHM of the LVC-NC components ranges from 12 to 39 km s⁻¹ with a median value of 27 km s⁻¹ and a standard deviation of 9 km s⁻¹. In contrast, the FWHM of the LVC-BC in this sample ranges from 49 to 140 km s⁻¹ with a median value of 97 km s⁻¹ and a standard deviation of 29 km s⁻¹. This cut off in FWHM between the LVC-NC and LVC-BC is adopted for this thesis to preserve consistency between this study, the subsequent studies following it in the literature (e.g. Banzatti et al., 2019; Fang et al., 2018) and the work presented in Chapters 4 and 5.

Fang et al. (2018) and Banzatti et al. (2019) expand on the work of Simon et al. (2016) to a further combined sample of over 100 stars in different star forming regions. Banzatti et al. (2019) find statistical correlations between the kinematics of the LVC (both velocity and FWHM) to the equivalent width of the HVC and the stellar accretion luminosity. Fang et al. (2018) find that the LVC-BC is more tightly correlated to the stellar accretion luminosity than the LVC-NC. They also observe that the ratio of [O I] $\lambda 5577$ to [O I] $\lambda 6300$ differs between the LVC-BC and the HVC, with the ratio being higher for the LVC-BC. The combined findings of these two studies reinforce the idea that the LVC-BC is tracing a hot/dense region at the base on an MHD disc wind. The origin of the LVC-NC as either an MHD (Banzatti et al., 2019; Fang et al., 2018) or thermal photoevaporative wind (Rab et al., 2023) currently remains in debate.

However, a key deficiency of these previous spectroscopic studies is the lack of spatial information needed to constrain the geometry of the wind. This is where the technique of as spectro-astrometry comes to the fore, providing this sought after spatial information on the LVC that previous spectroscopic studies lack. Chapter 4 uses high resolution spectroscopy (like these previous studies) in combination with the technique of spectro-astrometry to constrain the origin the LVC-NC and test the predictions of these wind driven accretion models by comparing the wind mass loss rates to the stellar mass accretion rates.

1.3 Branches of Brown Dwarf formation

The previous sections of this chapter have discussed the primary accretion and ejection mechanisms present in forming low-mass YSOs. In the upcoming sections, this discussion is briefly expanded to the sub-stellar mass regime, focussing on the main "branches" of Brown Dwarf (BD) formation. BDs are the intermediary mass objects between low-mass stars and giant planets. They are principally different objects compared to stars due to their low mass ($M_{BD} < 0.075 M_{\odot}$) and their inability to maintain the consistent fusion of hydrogen (Chabrier and Lenoble, 2023).

Prior to this discussion, it is apposite to first introduce the concept of the stellar Initial Mass Function (IMF). The IMF is a number distribution that describes the number of stars of particular initial masses born into a MC. This probabilistic distribution represents mass spread for a population of stars (Kroupa et al., 2026). Initially, the IMF was expressed in the form of a power law with an index of ≈ 2.35 (see Equation 1.6; Salpeter, 1955), where dN is the number of stars with masses in the range m and $m + dm$.

$$dN = \xi(m) dm ; \xi(m) \propto m^{2.35} \quad (1.6)$$

Since then, several works have shown that this initially derived form of the IMF suffered from biases effecting star count (Hopkins, 2018; Kroupa et al., 2013). Much observational work (e.g. Cheng et al., 2023; Villaume et al., 2017) and theoretical work (e.g. Chabrier, 2003; Kroupa, 2001) has been done to understand the functional form of the IMF as it has far reaching implications to many areas of astrophysics, from this thesis' topic of star and BD formation, to explaining the matter cycle in galaxies (see the review of Kroupa et al., 2026, and references therein). These works have suggested that the stellar IMF within the Milky Way galaxy peaks between ~ 0.1 to $0.3 M_{\odot}$ (Chabrier, 2003), implying that the majority of the material available to make stars is used to make low-mass stars rather than high-mass ($\geq 8 M_{\odot}$) stars or BDs. However, the stellar IMF is known to vary with time (redshift) and metallicity (Bate,

2025; Li et al., 2023).

The two main pathways of BD formation are the star-like and planet-like scenarios, as summarised in Fig. 1.11. These two branches act as a rudimentary (non definitive, Grieves et al., 2017) explanation to the so-called "BD desert". This desert refers to the seeming lack of BD mass companion objects in close orbits around host stars (Campbell, Walker, and Yang, 1988; Feng et al., 2022) compared to objects on longer orbits (Stevenson et al., 2023). This may hint at two separate populations of BDs, one where BDs form like stars and another in which BDs form as planets (see Palau et al., 2024, and references therein).

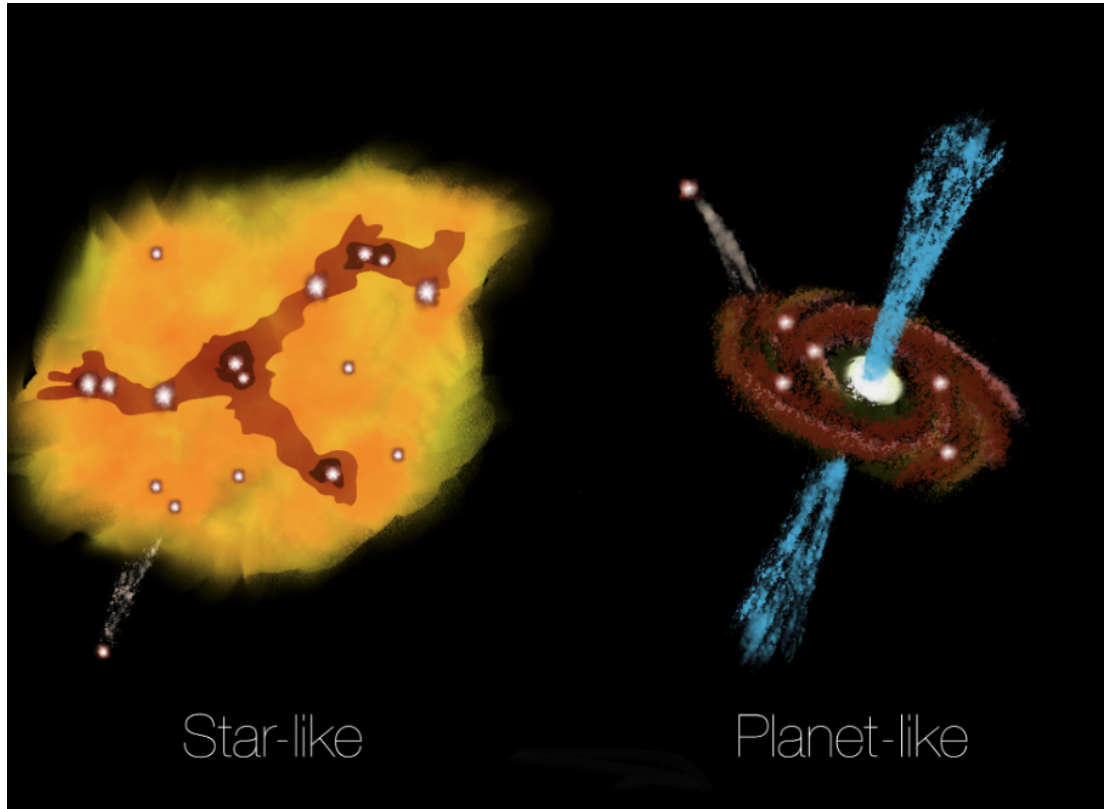


FIGURE 1.11: Branches of isolated brown dwarf formation, taken from Palau et al., 2024. In the star like scenario, the BD may be ejected from the molecular cloud during formation. In the planet-like scenario, the BD partly forms in-situ inside the circumstellar disc before subsequent ejection.

1.3.1 Star-like formation

In the star-like scenario, formation of the BD proceeds as a "scaled down" version of low-mass star formation where either the brown dwarf forms within turbulent fragments of molecular cloud cores (e.g. Bonnell, Clark, and Bate, 2008; Haugbølle, Padoan, and Nordlund, 2018; Padoan and Nordlund, 2002, 2004) or as a consequence of hierarchical gravitational collapse, whereby fragmentation can potentially occur isothermally down to masses as low as $\sim 0.005 M_{\odot}$ ($\sim 5 M_{Jupiter}$; see Larson, 1969; Young, 2023). A seemingly shared property of BDs that form in this way is a history of poor or stunted mass accretion. This may occur for several reasons, including but not limited to: the interaction of massive O or B type star stellar winds that strip away material available for accretion from the natal environment (Whitworth and Zinnecker, 2004; Whitworth, 2018), competition with YSOs for available material (Bonnell, 2005; Bonnell and Bate, 2006) or ejection from the birth environments (Bate, 2019; Reipurth and Mikkola, 2015). There is a series of observations that support the claim of the star-like scenario. For example, several studies of the IMF down to the sub-stellar mass regime in a few star forming regions have presented evidence indicating a log-normal functional form for the IMF at sub-stellar masses (e.g. Chabrier and Lenoble, 2023; Huston and Luhman, 2021; Kirkpatrick et al., 2019), which is consistent with what is expected for turbulence driven star formation (Zinnecker, 1984).

1.3.2 Planet-like formation

In the planet-like formation mechanism, BDs form or partially form within the confines of a circumstellar disc and are subsequently ejected through body-body interactions within the disc (Reipurth and Mikkola, 2015). Detections of low-mass free floating objects (e.g. Scholz et al., 2023) suggest some observational credibility for the planet-like formation mechanism. Additionally, a potential lower limit of $\sim 10 - 15 M_{Jupiter}$ has been suggested (Miret-Roig, N. et al., 2022; Schlaufman, 2018) as the minimum mass for star-like formation. What these previous studies of BDs lack is dedicated observations of the youngest and most embedded BDs the earliest stages

of formation. These are needed to distinguish between the "star-like" and "planet-like" formation scenarios for BDs and to determine the dominant formation pathway. Chapter 6 of this thesis provides dedicated molecular gas (emission line) observations of one such young and heavily embedded BD system using interferometry.

1.4 Variability

The formation processes described above are naturally variable in time, given that the star/BD at the centre of these systems and the surrounding discs evolves over time. The concluding section of this chapter now discuss some examples of variability observed for the accretion (Fischer et al., 2023) and ejection (Nony, T. et al., 2020) processes discussed above. The impact of this variability is obvious, in that it directly affects the central object, the surrounding disc and any possible forming planets. Although, there are still many questions in regard to the variability of these processes and this remains an active area of research. A particularly relevant example to this thesis is the variability observed in CTTs. The CTT DG Tau, shows variation in its photometric magnitude in time (Pyo et al., 2024), linked to changes in observed ejection patterns (Takami et al., 2023) and changes in accretion activity. DG Tau will be discussed in detail in Chapter 5.

1.4.1 Accretion variability

There are many types of accretion variability spanning many orders of magnitude and time as shown in Figure 1.12. A continuous accretion model fails to explain the observed under luminosity of stars (e.g. Kenyon et al., 1990) compared to what would be expected in the case of a simple continuous accretion model (for example, Kenyon and Hartmann, 1995). Accretion outbursts are part of the observed variable accretion activity and seem to offer a solution to this observed under luminosity problem (Dunham and Vorobyov, 2012). In oversimplified terms, accretion bursts in low-mass YSOs come in two main "flavours"; EX Lupi type outbursts and FU Ori type outbursts.

EX Lupi type bursts or EXors are characterised by an increase in photometric magnitude of up to ~ 2.5 mag and can last from several months to a few years (Sicilia-Aguilar et al., 2012). Their spectra resemble that of a CTTS, but they have strangely enhanced accretion (Fischer et al., 2023). EX Lupi itself has experienced three large outbursts, in the years 1945 (McLaughlin, 1946), 1955 (e.g. see Herbig, Suntzeff, and Blanco, 1992) and 2008 (Jones, 2008). A smaller outburst was detected in 2022 (Kóspál et al., 2022). The cause of these large outbursts in EX Lupi may be explained by clumpy accretion of material from the inner disc impacting the star thereby triggering the outburst phase (Sicilia-Aguilar et al., 2012).

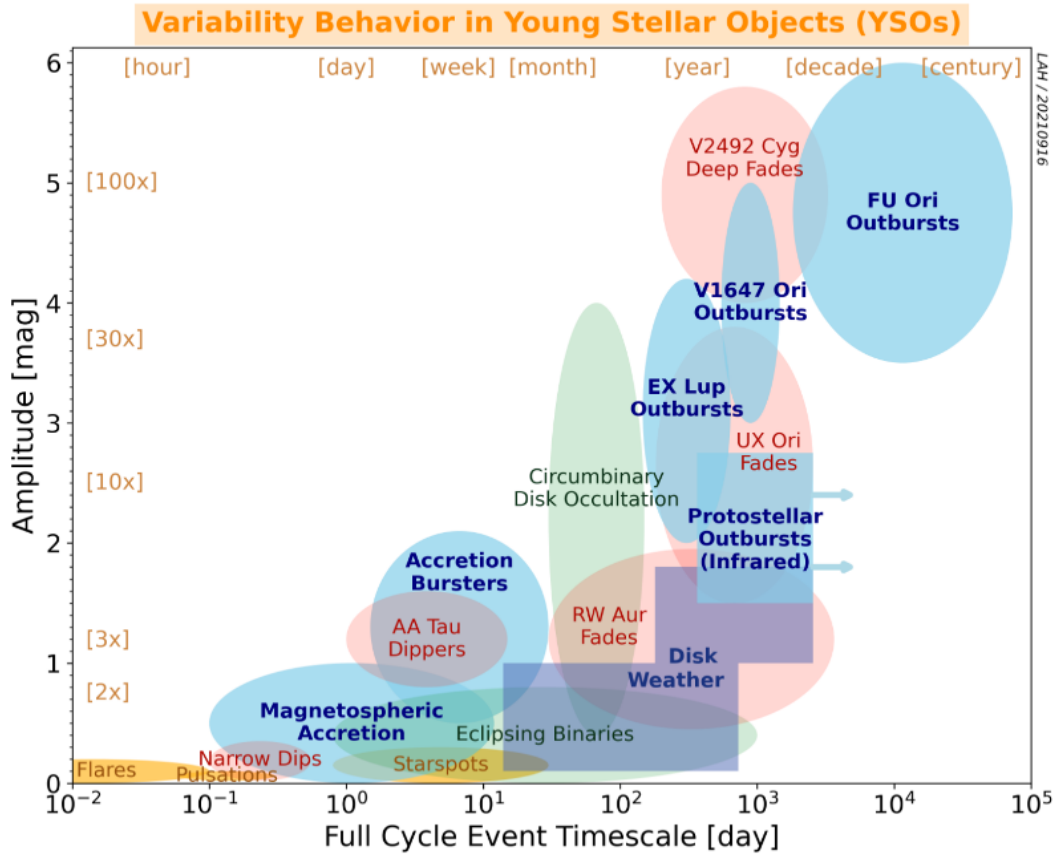


FIGURE 1.12: The many types of YSO accretion variability. These events span from hours to centuries in time and over six orders of magnitude. Figure taken from Fischer et al., 2023.

FU Ori like bursts or FUors, are objects which experience such violent outbursts that the flux from the disc outshines the central object (Liu et al., 2022). The outbursts span to photometric amplitudes ~ 4 to 6 mag and have decay times in the order of 100

years (Fischer et al., 2023). Their spectra are dominated by emission from the heated disc (see Liu et al., 2022, also).

One further example of accretion variability is the dipper subclass of YSOs. Dippers are YSOs that show periodic and/or irregular dips in their photometric light curves, corresponding to a significant proportion of the stars photospheric flux. A recent study (Empey et al., 2025) of 16 dipper objects using X-shooter spectra and Gaia photometry find 20 irregular dips that may be explained by the presence of dust structures in the inner disc obscuring the stellar photosphere. This explanation requires that the dust be somehow lifted off the disc via either unstable accretion columns (i.e. these columns vary in time during magnetospheric accretion e.g. Sousa et al., 2021) or disc winds (Miyake, Suzuki, and Inutsuka, 2016; Rodenkirch and Dullemond, 2022).

1.4.2 Outflow variability

In addition to accretion variability, YSOs also experience outflow variability. A key example of this is shown by the study of jet knots. Knots are internal shock structures in jets that are generated when fast, freshly ejected material (closer to the source) catches up to and slams into slower moving older material that is further above the ejecting source (e.g. Fig. 1.13).

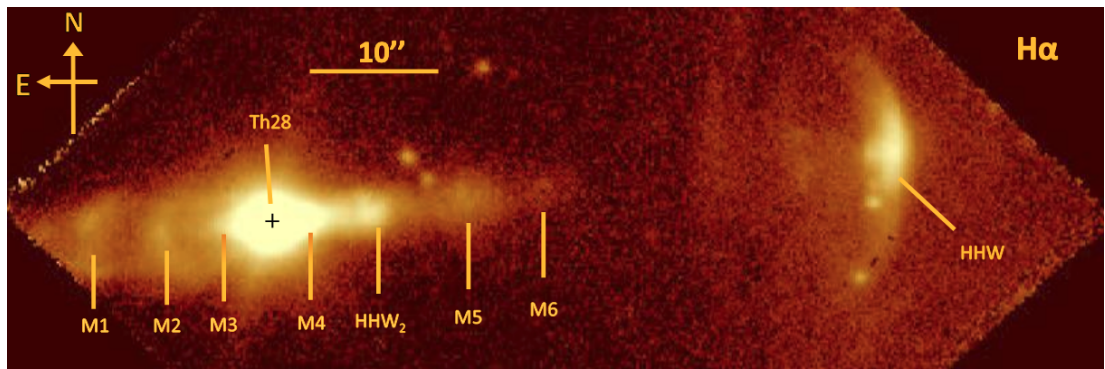


FIGURE 1.13: A spectro-image of the jet emitted from the CTTS Th28 in $H\alpha$ before continuum subtraction with the star (black cross), the jet knots and the bow shock marked. The image has been rotated such that the jet is along the horizontal axis. Taken from Murphy et al., 2021.

Knots in YSO jets are radiative and a significant amount of line emission can come from them (Ray and Ferreira, 2021). By examining the proper motions (a measure of the change in angular position as a function of time) of knots launched within jets, one can reconstruct the kinematic structure of the jet itself (Raga et al., 2017). Investigating the ejection timescale of these knots (e.g. Takami et al., 2023), and studying if knots have variable spacing between each other or changing proper motions could indicate a link between an ejected jet knot and a corresponding past accretion event (e.g. Raga et al., 1990; Vorobyov et al., 2018). Additionally, knots can be used to study the collision of jets in multiple systems (López et al., 2022). Given all of the above, it is clear that variability plays a significant role in the formation process of young stars and may be important in the similar star-like formation process of BDs.

Chapter 2

Spectroastrometry with the Ultraviolet & Visual Echelle Spectrograph

2.1 Spectro-astrometry: Its history, applications and limitations

This chapter begins with a description of the technique of spectro-astrometry (hereafter referred to as SA), an outline of the historical applications SA to various fields of astronomy from the techniques' inception to the present day, a discussion on its limitations and its practical application to the echelle observations taken with the Ultraviolet & Visual Echelle Spectrograph (UVES) presented in this thesis. The products of this technique underpin the results and discussion of the spatial properties of the optical FELs presented in Chapters 4 and 5. At its core SA is a technique which involves the fitting of the flux distribution (an emission line for example) to extract spatial information about that flux distribution, below what is available from diffraction limited observations of that flux distribution.

In practice, the flux distribution of an emission line region (ELR) in an echelle or long-slit spectrum is spread spatially across the slit by the observational seeing. Cuts taken of the flux distribution perpendicular to the dispersion direction at each row / column of pixels produces a profile of emission as a function of spatial spread across the slit. These profiles are referred to as "spatial profiles". Assuming the spatial profile is well sampled (at minimum Nyquist sampled) by the pixel scale of the detector; the spatial profile should be spread across multiple pixels of the detector and take on a characteristic Gaussian like shape in the first order approximation.

During SA, these spatial profiles are fit with a Gaussian function (Equation 2.1 where A , μ and σ represent the amplitude, mean and standard deviation from the mean respectively) to recover an estimate for the peak (referred to as the centroid) and FWHM of each spatial profile which covers the emission region of interest. The centroids are plot as a function of wavelength (or equivalently velocity if the rest wavelength is known) to produce what is known as a "position spectrum". A plot of the FWHM measurements versus velocity is typically plotted as ΔFWHM , where ΔFWHM is the individual FWHM measured for each spatial profile minus an average FWHM measured from a background/continuum region of the spectrum (Kirwan et al., 2022; Whelan et al., 2024). Generally, the average FWHM of ground based seeing limited observations is approximately equal to the seeing for that observation (Porter, Oudmaijer, and Baines, 2004). From hereafter, each plot of ΔFWHM versus velocity will be referred to as a " ΔFWHM spectrum".

$$G(x) = A \frac{1}{\sigma\sqrt{2\pi}} e^{-\frac{(x-\mu)^2}{2\sigma^2}} \quad (2.1)$$

Information gleamed from both the position spectrum and ΔFWHM spectrum can be used to investigate the spatial properties of an emission region with respect to the continuum emission. The SA measurements for this thesis are made as described above using a custom *Python* script. The steps that this script preforms are more deeply discussed in Section 2.1.2. In brief, the script extracts spatial profiles from the observed ELR, fits each spatial profile with a Gaussian profile and applies the necessary corrections and checks to construct the position and ΔFWHM spectra for each ELR of interest. The next section shortly outlines the history and historical and present applications of SA.

2.1.1 History of spectro-astrometry and its astrophysical applications

The first ideas leading to the development of SA came from the need to extract sub seeing information (i.e. to remove the effect of atmospheric turbulence on images).

These first ideas were recorded in the early 1980s in the proceedings of the International Astronomical Union (IAU) Colloquium no. 62 "Current techniques in double and multiple star research". Hence, SA traces its founding back to earlier techniques such as chromatic position difference (CPD) and differential speckle interferometry (DSI) which were used to separate unresolved binary stars (e.g. Beavers et al., 1989; Beckers, 1982; Christy, Wellnitz, and Currie, 1981). These early techniques were refined over time and the term SA was first coined by Bailey (1998).

Since then, SA has proved to be a powerful technique in many areas of astrophysics. Initially used to detect unresolved spectroscopic binary stars (Bailey, 1998), SA has been used for a variety of purposes. Some examples include; the characterisation of the gas in discs around Herbig stars (Brittain, Najita, and Carr, 2015), a possible exoplanet detection technique (Cahill et al., 2019), measuring the geometric distances to extragalactic binary systems (Songsheng et al., 2025), studying possibly asymmetric structures in the environment around asymptotic giant branch stars (Hron et al., 2015), to search for discs inside planetary nebulae (Blanco Cárdenas et al., 2014), to study the close in discs and envelopes around massive YSOs (Grave and Kumar, 2007) and to examine the rotating gas discs around active galactic nuclei as a means constrain their kinematics and to estimate the mass of the central black hole (Gnerucci et al., 2010, 2011, 2013).

For the purposes of this thesis, SA is used to study the outflows launched from low-mass YSOs during their formation process. The study of outflows with SA has been ongoing for more than twenty years. The technique has examined outflows from both loss-mass YSOs. (e.g. Hirth, Mundt, and Solf, 1997; Solf, 1984; Takami et al., 2001; Whelan, Murphy, and Pascucci, 2023; Whelan et al., 2024) and BDs (e.g. Whelan et al., 2014; Whelan et al., 2005) in both atomic and molecular emission lines (e.g. Davis et al., 2003; Davis et al., 2001; Pontoppidan, Blake, and Smette, 2011). So, it is clear since its inception that SA has been a versatile and adaptable technique suitable to address many areas of research where the observational seeing limits the spatial information

available to more traditional diffraction limited techniques. The next section discusses in detail, the practical application of SA to echelle and/or long-slit spectra and its current limitations.

2.1.2 Performing spectro-astrometry echelle/long-slit spectra

The first step in performing SA is obtaining suitable and clean data. The most suitable data is intermediate spectral resolution high signal-to-noise ratio (S/N) spectra (Whe- lan and Garcia, 2008) that is free from the contamination of cosmic ray effected pixels, skylines etc (see Sections 2.2.2 and 2.2.3 for a discussion on the post pipeline reduction steps done on the UVES data). Another key consideration is the position angle (PA) at which each spectrum is taken. Figure 2.1 summarises a typical observation setup for the required slit PAs needed to perform SA on a stellar system that drives a bipolar outflow. This example was chosen as it is the most relevant example given the work presented in Chapters 4 and 5.

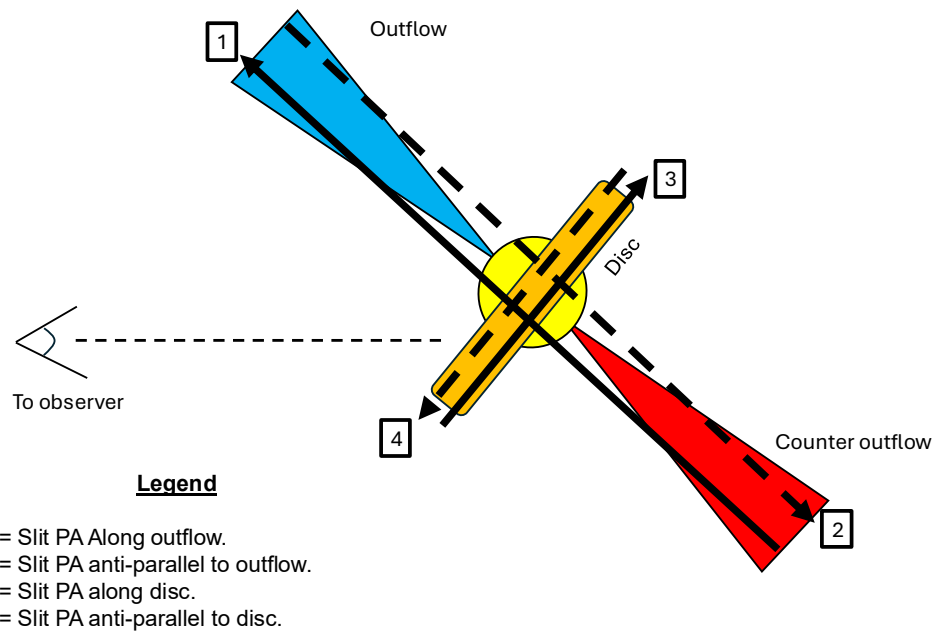


FIGURE 2.1: Schematic diagram of the slit PAs used for the spectro-astrometric technique.

Observations at four slit positions are taken, one along the outflow axis, one anti-parallel to the outflow axis, one along the disc major axis and one anti-parallel to disc

major axis in 90° increments. The orientation of the slit PA to the actual PA of the extended emission (in this example a jet) should be as close as possible so that any measured spectro-astrometric offset will be maximised. The anti-parallel slit PAs are taken as this is a typical strategy used to account for the effect of spatial artefacts on the position spectrum (Brannigan et al., 2006). The perpendicular slit PAs are taken so that a 2-dimensional (2D) map of the emission region can be constructed (Whelan et al., 2012). Spatial artefacts are false centroid offsets in the position spectrum that can be caused for a few reasons (e.g. poor tracking or flat fielding) but, they can be accounted for by the subtraction of centroid offsets measured in an anti-parallel observation. A schematic diagram of this process is presented in Fig 2.2 and summarised in Equation 2.2. In this equation, $x_{c(true)}$ is the true centroid offset, $x_{c(0^\circ)}$ is the centroid offset along one slit PA and $x_{c(180^\circ)}$ is the centroid offset along the corresponding anti-parallel slit PA. The *Python* script developed for the purpose of preforming SA for this thesis takes this effect into account and calculates the true centroid offset.

$$x_{c(true)} = \frac{1}{2} \times (x_{c(0^\circ)} - x_{c(180^\circ)}) \quad (2.2)$$

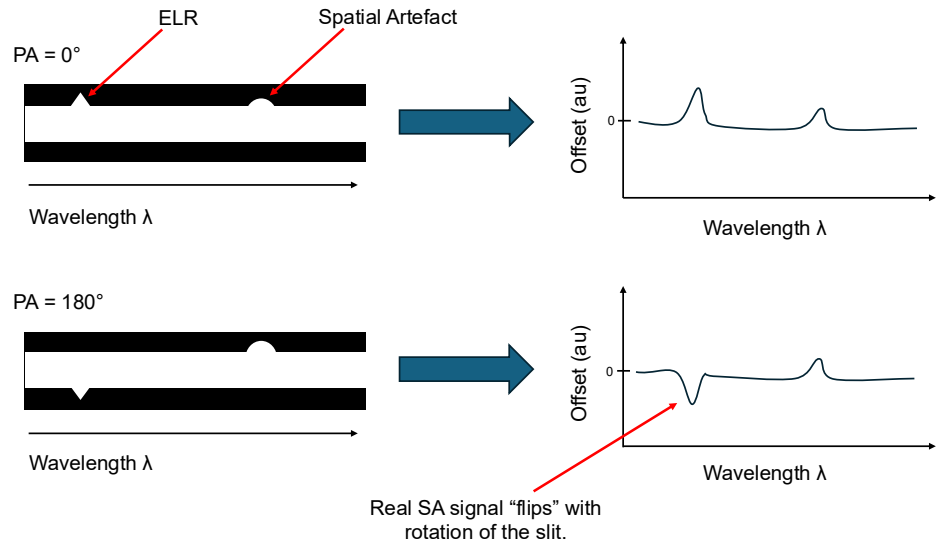


FIGURE 2.2: Schematic diagram showing the procedure for the removal of time independent spectro-astrometric artefacts. Left: Cartoon 2D spectra at two anti-parallel slit PAs. Right: An exaggerated position spectrum of each PA.

The schematic presented in Fig.2.2 shows that any real shift in the position spectrum due to an ELR should "flip" as the slit is rotated through 180° while any remaining false shift due to a spatial artefact remains in place as the slit is rotated. Next is a description of how the position and ΔFWHM spectra measured in practice. This was briefly discussed above, but there are several other effects alongside spatial artefacts that need to be considered to answer this question. Figure. 2.3, presents an example of a fit to a one spatial profile (done using the custom *Python* script) of the [O I] $\lambda 6300$ emission line of the CTTS DG Tau before and after the removal of the background continuum emission from the star. The effect of the continuum has been removed as the measured centroid in each spatial profile is weighted by the flux of each contributor to the overall flux distribution. In more simple language, the contribution of the continuum compared to the contribution of an ELR to a flux distribution is often large. Given this situation, any centroid measured will be weighted towards the position of the continuum source (e.g. a star). So, in order for the true positional offset of the pure ELR to be recovered, this contaminating effect of the continuum emission on the centroid measurements has to be removed.

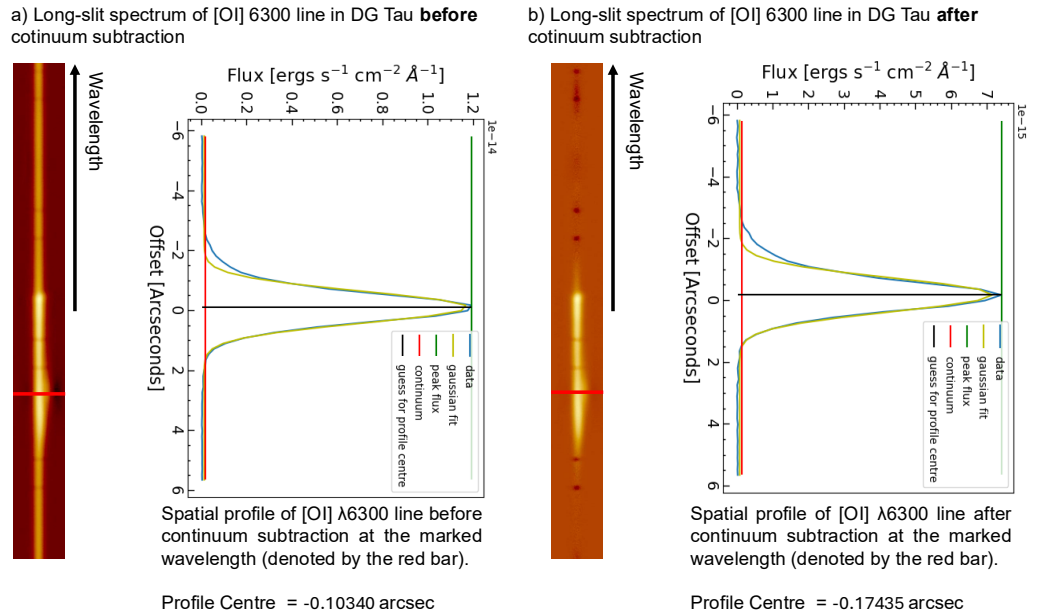


FIGURE 2.3: Example of the changes in the measured spectroastrometric centroid for the [O I] $\lambda 6300$ line in DG Tau. a) before and b) after continuum subtraction, hence leaving the offset due only to the pure ELR.

This can be done in two ways. Firstly, as demonstrated in Fig. 2.3, the peak and FWHM of each spatial profile in the ELR can be measured before and after the careful subtraction of the continuum emission from the 2D spectrum (as detailed in Section 2.2.3). A second approach was introduced by Takami et al. (2001) in which the effect of the continuum emission is accounted for by calculating a correction factor at each wavelength/velocity. This correction factor is given in Equation 2.3 and is the ratio of the ELR flux to the ELR flux plus the continuum flux. Once this factor is calculated, the true spectro-astrometric offset for the ELR may be found by dividing the uncorrected spectro-astrometric offsets calculated prior to the continuum subtraction by this factor.

$$I_{\lambda(line)} / (I_{\lambda(line)} + I_{\lambda(cont)}) \quad (2.3)$$

Both of the methods that account for the effect of the continuum have their advantages and disadvantages. For example, the continuum subtraction method has the advantage of allowing for the extraction of spectro-astrometric signal from ELRs that may be "buried" below the continuum emission. A disadvantage to this method is that it requires careful and precise fitting and removal of the continuum emission. It also requires that each spatial profile is fit both before and after continuum subtraction, adding some extra work. The correction factor method has the advantage of avoiding the need to fit the spatial profiles twice. However, a key disadvantage of this method is that it only amplifies existing spectro-astrometric signal from the ELR found prior to the continuum removal (i.e. any small spectro-astrometric offset due to part of an ELR that may not be detectable prior to the removal of the background continuum emission could possibly be missed when the spectro-astrometric offset is scaled by the correction factor shown in Equation 2.3).

In principle, both methods should produce similar results however, in this thesis, the continuum subtraction method is used. This method is preferred over the correction factor method for the advantage of allowing for the extraction of small spectro-astrometric signals from ELRs that may be "buried" below the continuum emission. This could be a distinct possibility given that the optical FEL LVC studied in Chapters

4 and 5 may be very spatially compact (Fang et al., 2023b). Figure 2.4 presents an example made in *Python*, showing the effect of the continuum on position spectrum of the [O I] λ 6300 FEL region of the CTTS DG Tau along the axis of the stars bipolar jet before and after the continuum subtraction routine. Additionally, Table 2.1 presents a comparison of the pixel size to the seeing for all the observations presented in Chapters 4 and 5.

TABLE 2.1: Comparison of the seeing to pixel scale of all observations presented in Chapters 4 and 5.

Object	Observation year	Seeing (arcsec)	Pixel scale (arcsec per pixel)
TW Hya	2012	0".53	0.182
T Cha	2012	0".90	0.182
MP Mus	2012	0".64	0.182
LkH α 330	2021	1".22	0.182
GQ Lup	2022	1".28	0.182
CW Tau	2021	0".88	0.182
GK Tau	2021	1".06	0.182
DK Tau	2022	1".03	0.182
EX Lup	2022	1".10	0.182
DG Tau	2003	0".26	0.138 ^a
DG Tau	2010	1".16	0.182
DG Tau	2021	1".04	0.182

^a Observation taken with the High Dispersion Spectrograph (HDS) on the SUBARU telescope.

One more consideration is to ensure that the position spectra and Δ FWHM spectra are measured relative to the position of the continuum source (i.e. the continuum source is positioned at an offset of 0 arcseconds or 0 astronomical units (AU)). However, any residual tilt or curvature in the 2D echelle or long-slit spectrum used for the SA analysis would introduce a "shift" in the position of the continuum source (Whelan and Garcia, 2008). To correct for this, a polynomial is fit to the position spectrum (in the regions free of line emission) prior to the continuum subtraction (i.e. the position spectrum is still contaminated by the continuum). This polynomial is then subtracted from the position spectrum **of both** the continuum and non continuum subtracted centroids, to move the position of the continuum source back into the zero offset position. This process is also part of the custom *Python* script used to preform SA on the data presented in this thesis and is illustrated by Fig. 2.5.

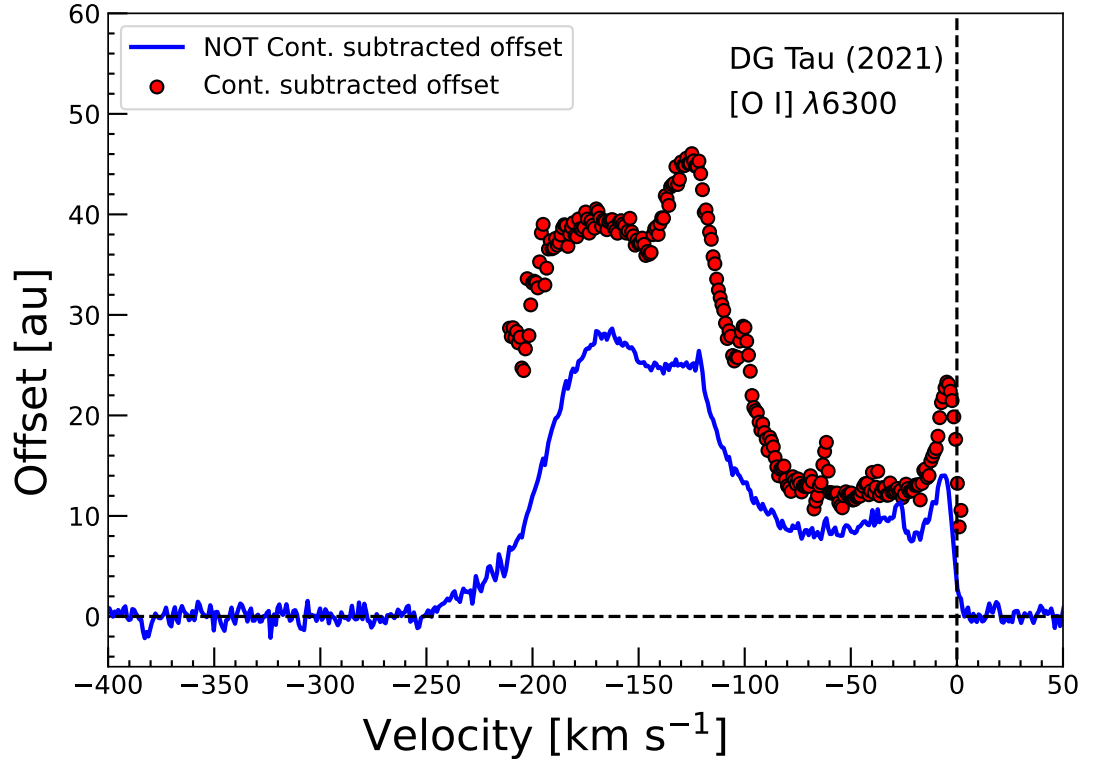


FIGURE 2.4: Example showing the effect on the measured position spectrum before and after continuum subtraction for the [O I] $\lambda 6300$ line of DG Tau along the outflow axis ($PA = 46^\circ$) in the 2021 UVES data. Zero offset represents the position of the star. The velocity axis is in the stellocentric rest frame.

With the above information, spectro-astrometric analysis provides valuable insights into ELRs. The position spectrum gives the observer a lower limit estimate on the position of an ELR (weighted by the brightest emission in the ELR) with respect to the position of the continuum source whereas the $\Delta FWHM$ spectrum provides the observer with an estimate of the intrinsic size of the emitting region. Section 2.1.3 expands on the discussion above by examining the accuracy and limitations of SA.

2.1.3 The accuracy and limitations of spectro-astrometry

The accuracy to which any given centroid measurement in the position spectrum can be measured to is given by Equation 2.4 (Brittain, Najita, and Carr, 2015; Whelan and Garcia, 2008)

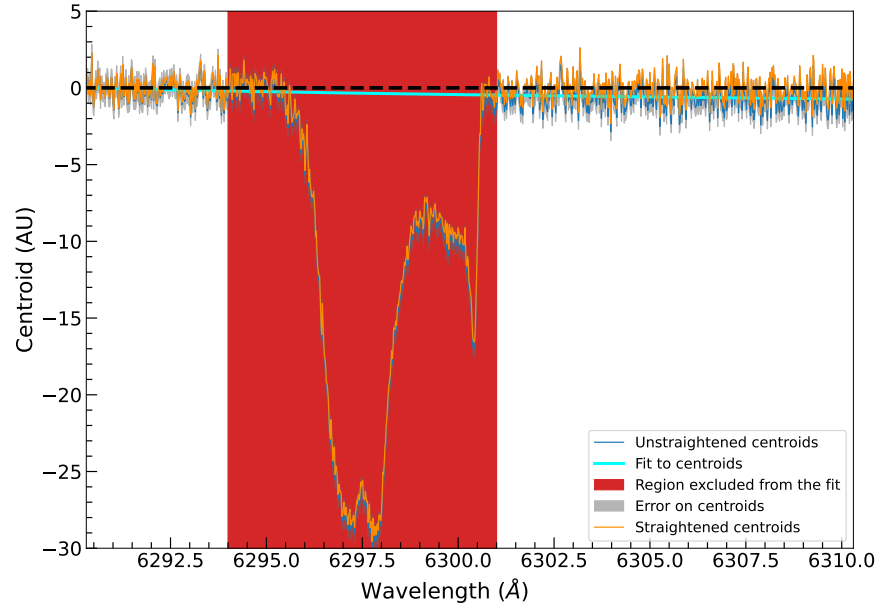


FIGURE 2.5: Example of a polynomial fit to the continuum contaminated position spectrum. This accounts for residual tilt/curvature in 2D spectrum. The black dashed line corresponds to zero offset. This should be the position of the continuum region centroids after the straightening procedure has been performed. The centroids associated to the ELR have been excluded from the straightening fit (marked in red).

$$\sigma_{centroid} = \frac{seeing \text{ (in mas)}}{2.3548 \times \sqrt{N_p}} \approx \frac{0.4 \times FWHM}{S/N} \quad (2.4)$$

In this equation, N_p is the number of detected photons and the FWHM of the spatial profile can be used to estimate the seeing (Porter, Oudmaijer, and Baines, 2004). The value of N_p may be estimated from the counts of the spatial profile (i.e. the area under the curve of the spatial profile before flux calibration) multiplied by the gain of the detector. The uncertainty on the centroid measurement can be improved by applying a running average to the spectrum (also known as binning) which increases the S/N at the expense of degrading the spectral resolution and also the extent of the measured offsets (by averaging the peaks of adjacent spatial profiles). In each of the position spectra presented in Chapters 4 and 5, the quoted uncertainty on each position spectrum is an average of the values calculated for each spatial profile using Equation 2.4.

The uncertainty on each individual spatial profiles' FWHM is determined from the Gaussian fit. The `curve_fit` function in the *Scipy* module (Virtanen et al., 2020) used in

Python to fit a Gaussian function to each spatial profile produces a so-called "covariance matrix" for each fit. This matrix contains all zeros except for the diagonal entries (in the assumption that no parameters of the Gaussian fit are dependent on one another). These diagonal entries are the variances on the parameters of the Gaussian function used to fit each spatial profile. Taking the square root of these variances gives the standard deviation or uncertainty on each parameter of the Gaussian fit. Given this, the uncertainty on the FWHM of each spatial profile is then defined as in Equation 2.5. In this equation σ represents the standard deviation (or fit uncertainty) on the fit parameter σ in Equation 2.1.

$$\sigma_{FWHM} = 2\sqrt{2 \ln 2} \times \sqrt{\sigma^2} \quad (2.5)$$

The uncertainty on the $\Delta FWHM$ values of the ELR of the spectrum is taken to be the average value of σ_{FWHM} in the continuum region of the spectrum. This value is the uncertainty quoted on each $\Delta FWHM$ spectrum presented in Chapters 4 and 5.

There are two main limitations to SA. The first is that SA only provides a lower limit on the extent of ELR. This lower limit estimate is weighted by the brightest emission in the ELR. To obtain the full/true extent of the emission traced by an ELR, techniques such as spectro-imaging may be required (Birney et al., 2024a; Fang et al., 2023b; Kirwan et al., 2022; Murphy et al., 2021, 2024). However, the advantage of SA over spectro-imaging is the ability to probe the extended emission very close to its emitting region.

The second limitation of SA is the need for asymmetric structure to obtain a signal. As illustrated in Fig. 2.6, if SA is performed on some extended structure with symmetric geometry (e.g. some spherical shell of emission), the offset measured in the position spectrum is observed at the same pixel with in both directions at once. So, the SA signal effectively "cancels itself out".

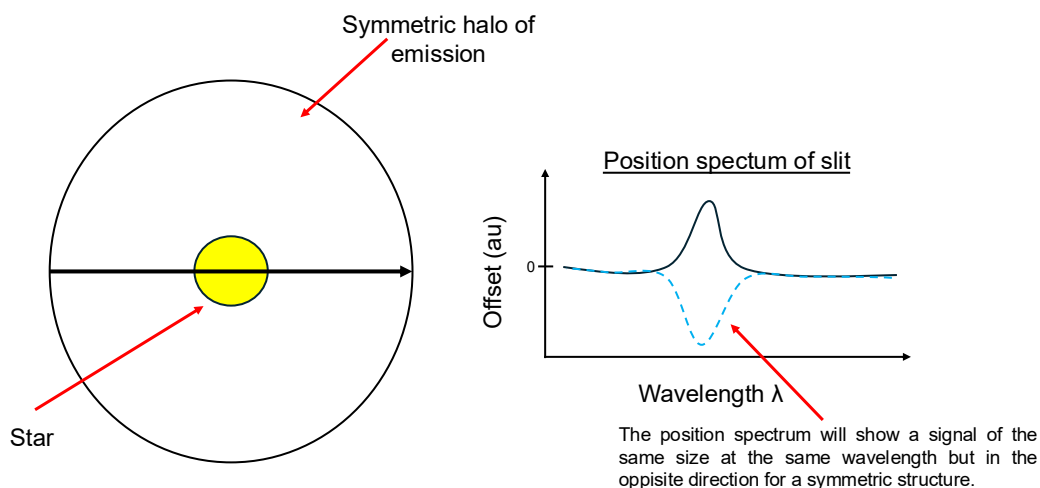


FIGURE 2.6: Cartoon showing how the spectro-astrometric signal in a symmetric structure cancels itself out as signal is detected in both directions simultaneously at the same wavelength.

2.2 The Ultraviolet and Visual Echelle Spectrograph

The remainder of this chapter will discuss the UVES instrument, data reduction pipeline and post processing steps performed after the initial pipeline reduction. The UVES instrument was the primary tool used to collect the spectra for the spectro-astrometric analysis presented in this thesis. This instrument was selected as it is well known that UVES is "work horse" instrument for doing SA (Whelan, Murphy, and Pascucci, 2023; Whelan et al., 2021; Whelan et al., 2024). X-Shooter (Vernet et al., 2011), has lower spectral resolution ($R = 6700$ for a $1''.2$ slit in the VIS wavelength range of the instrument (559.5 to 1024 nm); Vernet et al., 2011) compared to UVES in addition to documented issues with spatial artefacts. X-Shooter suffers from these spatial artefacts potentially due to the re-binning of the wavelength axis during the pipeline data reduction procedure as described by Whelan et al., 2015. Given this, UVES was selected over X-Shooter for the collection of data suitable for spectro-astrometric analysis.

UVES (Dekker et al., 2000) is a high spectral resolution cross-dispersed optical spectrograph located at the Nasmyth B focus of Unit Telescope 2 (UT2), of the ESOs' Very Large Telescope (VLT). UVES consists of two arms. A UV to Blue arm and Visual to Red arm covering a combined wavelength range from ~ 300 nm to ~ 1100 nm. The

spectrograph has a maximum resolving power of $R \sim 80,000$ and $110,000$ in the Blue and Red arms respectively. The instrument contains three $2k \times 4k$ charged coupled detectors (CCDs), (two detectors for the Red arm with different mosaic types) and a two other CCDs for slit acquisition. Figure 2.7 presents a schematic view of the instrument.

2.2.1 The UVES data reduction pipeline

After acquiring the unprocessed UVES data (science frames, lamp frames, etc.), these data were initially reduced using the ESOs' UVES data reduction pipeline (Freudling et al., 2013a) within the *ESO Reflex* environment. An overview of the pipeline interface and reduction steps is presented in Fig. 2.8. Below is a brief outline the initial data reduction steps performed by the UVES pipeline.

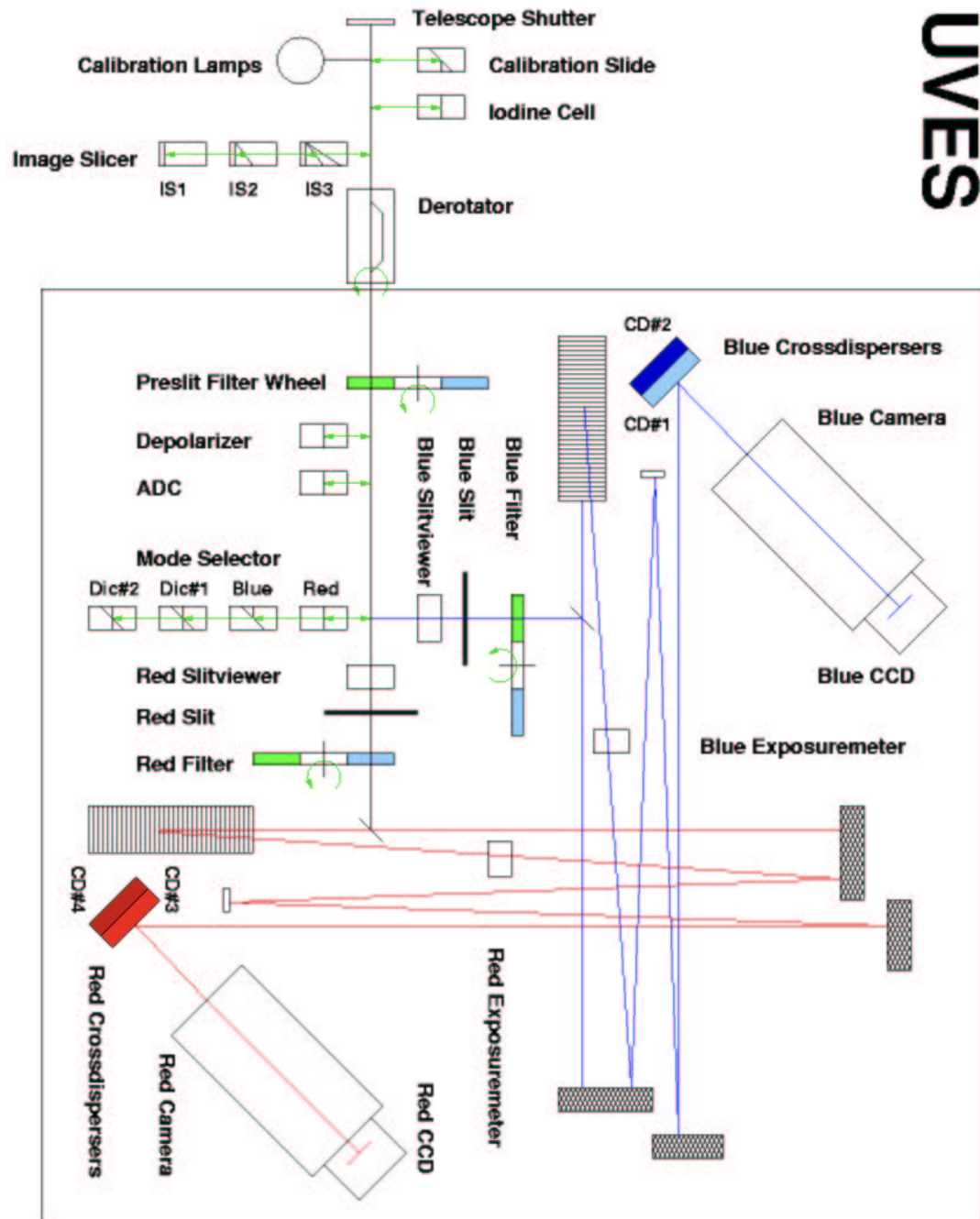


FIGURE 2.7: Schematic overview of the UVES spectrograph. This image was reproduced from the UVES user manual version 115 (available at https://www.eso.org/sci/facilities/paranal/instruments/uves/doc/ESO_514367_User_Manual_P115.pdf) with permission from L. Sbordone.

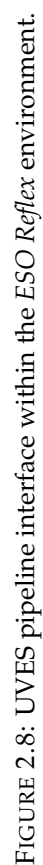


FIGURE 2.8: UVES pipeline interface within the *ESO Reflex* environment.

The pipeline begins by organising the unprocessed data into datasets which may contain one or more science files together with all the calibration files needed to process that dataset. The files that belong to each dataset and the purpose of each file during the reduction process is determined by OCA¹ rules. Once the data is organised, the pipeline generates a master bias, dark and flat frame via median combination of corresponding frames in the dataset. In this second step, the pipeline also detects the positions of the echelle orders and makes an initial guess of the wavelength calibration solution to be refined in the third step. In the third step the pipeline refines the wavelength calibration and performs the flux calibration. The wavelength calibration is done by comparing the initial guess of wavelengths of the strongest lines observed in the lamp frame (done in the second step) to a table of known reference wavelength values that come pre-packaged with the *ESO Reflex* workflow.

The flux calibration is done by first reducing the spectrum of a flux standard star (i.e. a star with known/well measured flux) in the exact same way as the science data (i.e. using the same master calibration files). The pipeline then uses the flux standard star to find a precalculated physical flux value for that standard in a flux reference table provided within the pipeline package. The flux standard star spectrum is divided by this flux and corrected for several effects (e.g. exposure time, detector binning etc). From this, the pipeline creates a response curve which is used to flux calibrate the science spectra. An example of an instrument response curve with the associated non-flux calibrated science spectrum is presented in Fig. 2.9. The flux calibration is skipped if no flux standard star spectrum is provided to the pipeline.

¹OCA stands for Organisation Classification and Association. These are the rules which classify data files according to the contents of each file header and sets how files should be grouped for processing by the pipeline.

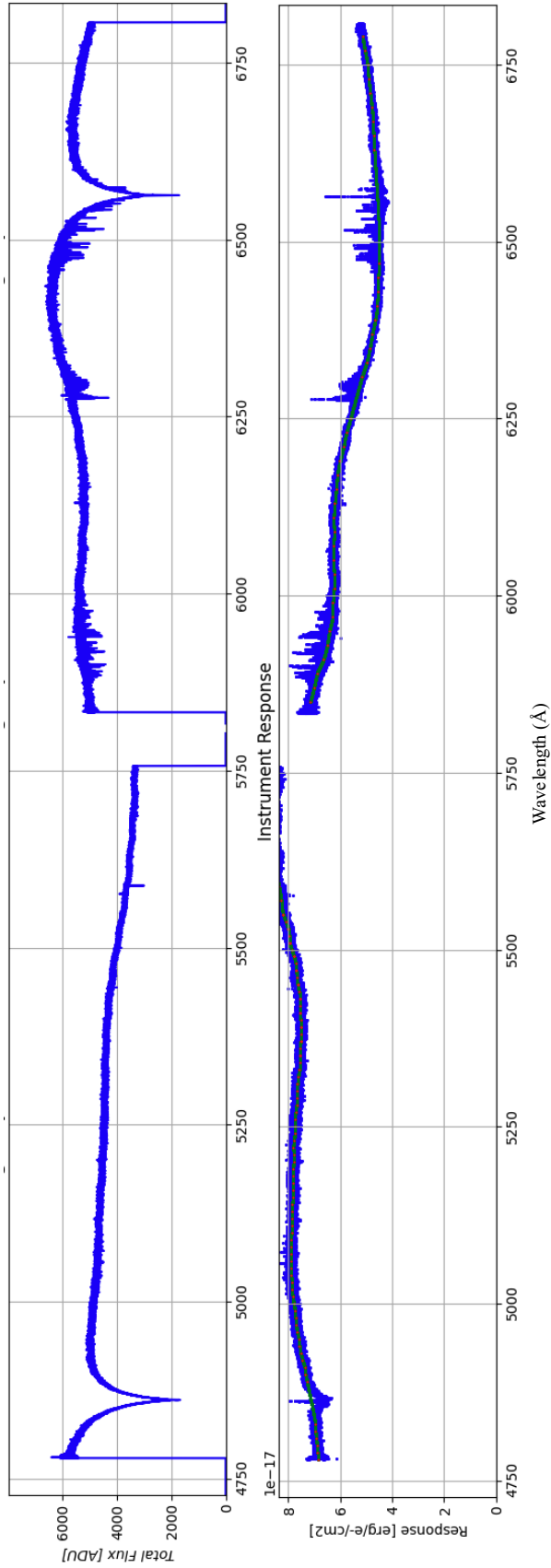


FIGURE 2.9: Top panel: Example on science spectrum before flux calibration. Bottom panel: The corresponding instrument response curve used to perform flux calibration.

The last major step of this initial reduction is the extraction of the calibrated spectra from the echelle data. The UVES pipeline allows the user to modify the length, width and offset (from the centre of the echelle order) of the spectrum extraction "window" for optimization in various use cases (e.g. point source versus extended source). Other "windows" may be used during the spectrum extraction process depending on the extraction method selected. For example, "windows" sampling the background may be used for sky subtraction. Figure 2.10 presents a schematic diagram outlining the spectrum extraction process. Two different extraction methods are used to extract 1-dimensional (1D) and 2D spectra "optimal" and "2D" respectively. For 1D spectra; the flux within the extraction slit is summed. For 2D spectra, the extraction slit is resampled into an image with wavelength on the x-axis and position along the slit on the y-axis.

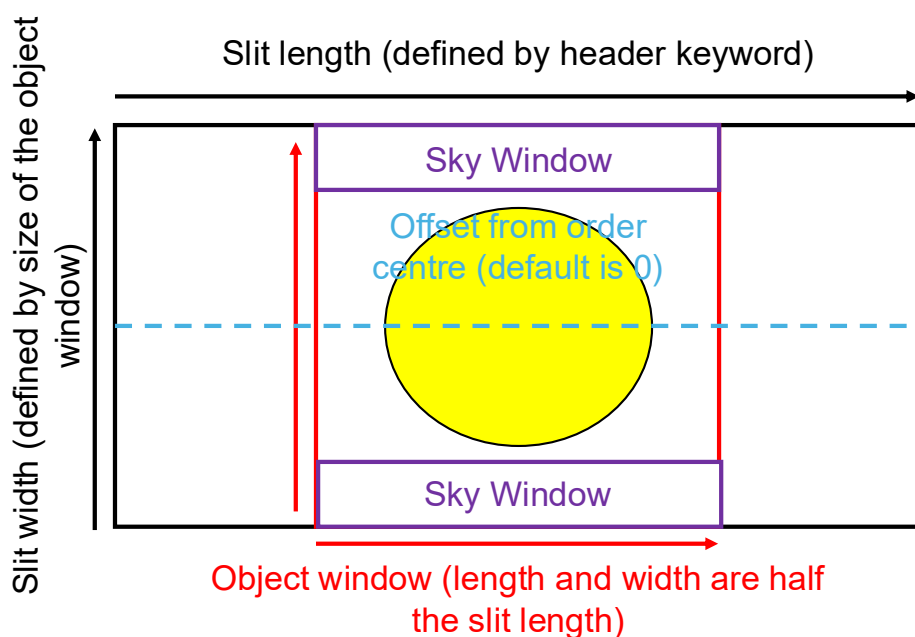


FIGURE 2.10: Schematic of the spectrum extraction process in the UVES pipeline. The coloured boxes show the various regions defined during spectrum extraction via the UVES pipeline workflow recipe `uves_obs_scired`. The point source (in yellow) is centred in the slit for visualisation purposes.

2.2.2 Post pipeline data processing: 1D spectra

Following the initial data reduction and calibration of 1D and 2D spectra obtained for each CTTS from the UVES pipeline, several post processing steps were done to both the 1D and 2D spectra to make them suitable for scientific analysis. The following subsection provides a discussion on the post processing of the 1D spectra.

Telluric and photospheric corrections

The initial 1D spectra that emerge from the UVES pipeline remain contaminated by two types of unwanted absorption features. Telluric absorption appears as a series of deep and narrow features imprinted on the target spectrum when some of the targets light is absorbed by molecules (e.g. O_2 , H_2O) in Earths' atmosphere. Photospheric absorption are broad and shallow features that also contaminate the target spectrum. These features are generated when certain wavelengths of light produced at the stellar surface are absorbed by opaque layers of the stellar atmosphere before escaping into the ISM and travelling toward the observer. The stellar rotation impacts the observed width of these features. Veiling due to stellar mass accretion can "fill in" the photospheric features in the science spectrum making them appear to be more shallow due to the excess flux generated during stellar accretion (Calvet and Gullbring, 1998; Stock et al., 2022). Figure 2.11 presents an example of an $[O\text{I}] \lambda 6300$ line from a CTTS contaminated by telluric and photospheric absorption.

There are two strategies typically used to remove telluric contamination from the science spectrum. The first is to observe a telluric standard star (i.e. a hot and fast rotating O or B type star with no intrinsic features) with the same instrument close in time to the science observation. These standard star observations are normalised to their continuum level and divided into the science spectrum to remove the telluric features. The advantage of this method is that it generates a spectrum of the telluric features that is at the same spectral resolution and convolved with the same instrument response as the science spectrum. A disadvantage of this method requires dedicated

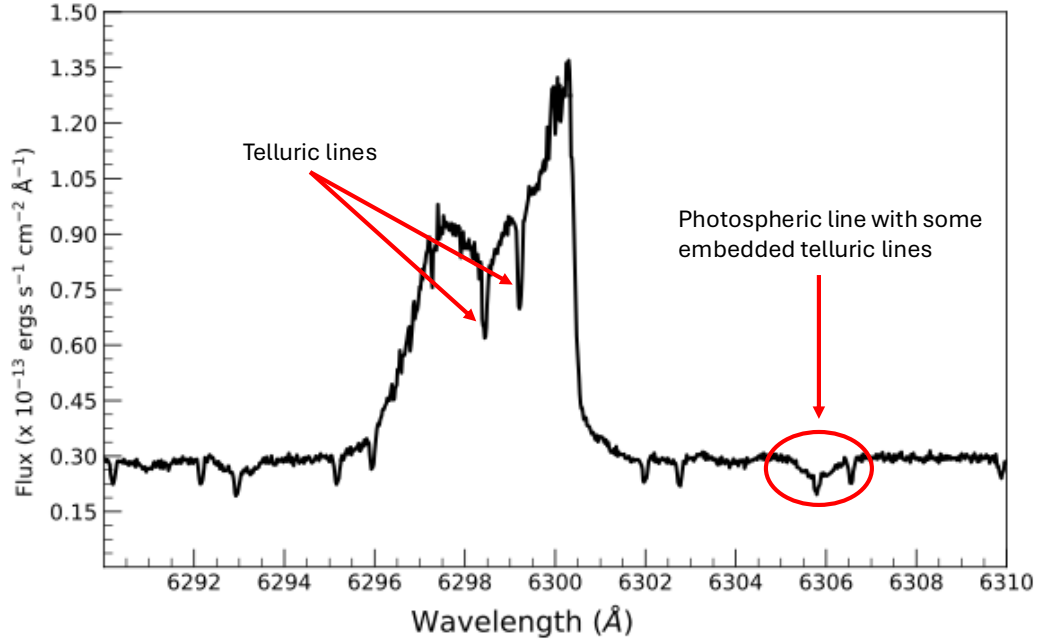


FIGURE 2.11: Example of a [O I] $\lambda 6300$ line in the telescope rest frame after processing and extraction from the UVES pipeline with telluric and photospheric absorption lines present.

observations of the telluric standard star close in time to the science target observations. A large difference in time between the target and telluric standard observations introduces changes in the atmospheric conditions which may produce spectral artefacts in the telluric subtracted target spectrum.

The second method to remove telluric absorption features is to create a model of the atmospheric transmission at the time that the science spectrum is observed. This is the method chosen for this thesis as telluric standard observations were not taken during science observations. The ESOs' telluric correction pipeline, MOLECFIT (Kausch et al., 2015; Smette et al., 2015) uses this second approach to remove telluric contamination from the 1D science spectra. The synthetic transmission approach has been shown to be an effective tool for removing telluric absorption up to the near infrared (Ulmer-Moll et al., 2019). An overview of the MOLECFIT user interface is shown in Fig 2.12.

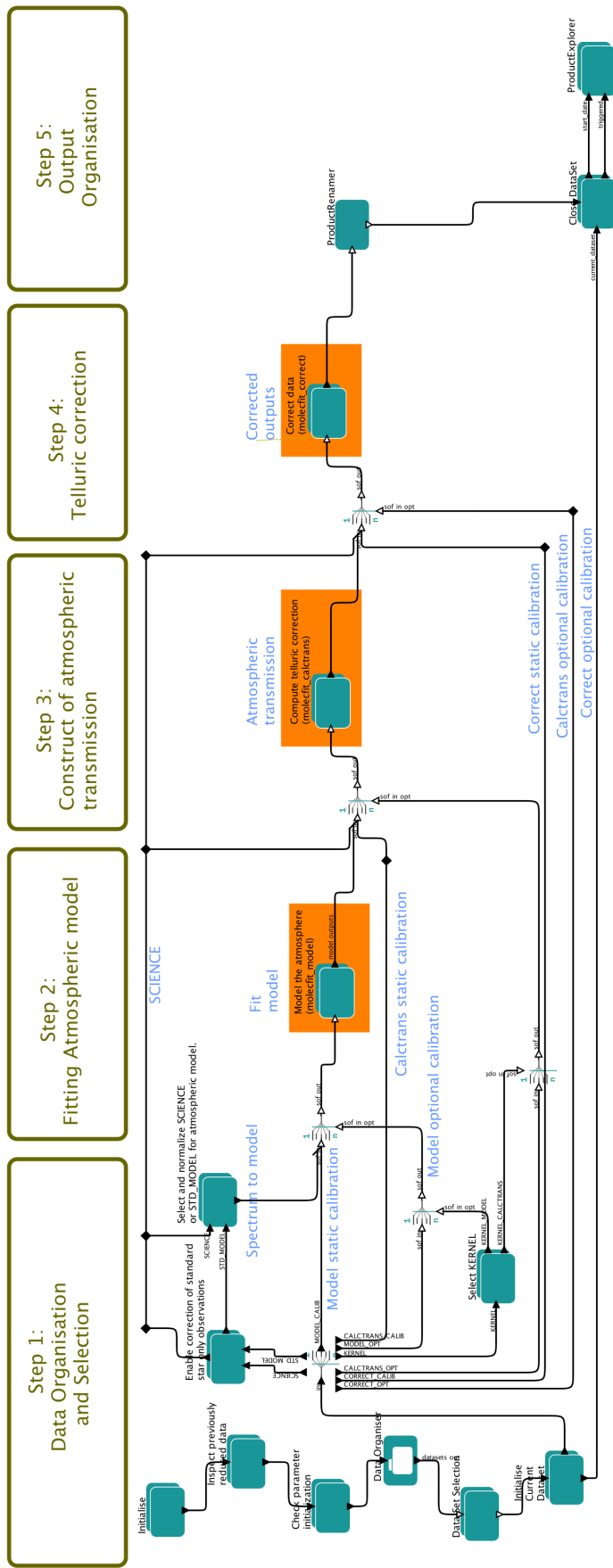


FIGURE 2.12: MOLECFIT telluric correction pipeline interface within the ESO Reflex environment.

MOLECFIT uses metrological measurements (taken at the observatory at the time of the target observations), the column densities of the molecules that shape the telluric lines (e.g. CO₂, O₂ etc.) as a function of altitude and location in combination with a radiative transfer code to construct a model of the atmospheric transmission at the time of the science observation. The telluric contamination is then removed by dividing the model into the science spectrum.

It is noted here that the column densities used by MOLECFIT are derived by measuring specific absorption features in a reference spectrum. This reference spectrum can either be an observation of a telluric standard star taken at any time or the science target itself if it is bright enough. An advantage to this second method is that dedicated observations of telluric standards will be minimised, freeing up more time for science observations. A disadvantage of this method is that it relies on how the transmission model treats the physical properties of each molecule that shapes the telluric lines and how these molecules interact with each other. This model also assumes that the modelled variations in atmospheric conditions with time is accurate enough to produce the correct absorption line strengths. An example of the telluric correction is shown in Fig. 2.13. A detailed description of the MOLECFIT pipeline is given in the MOLECFIT user manual available at <https://ftp.eso.org/pub/dfs/pipelines/instruments/molecfit/molecfit-pipeline-manual-4.4.2.pdf>.

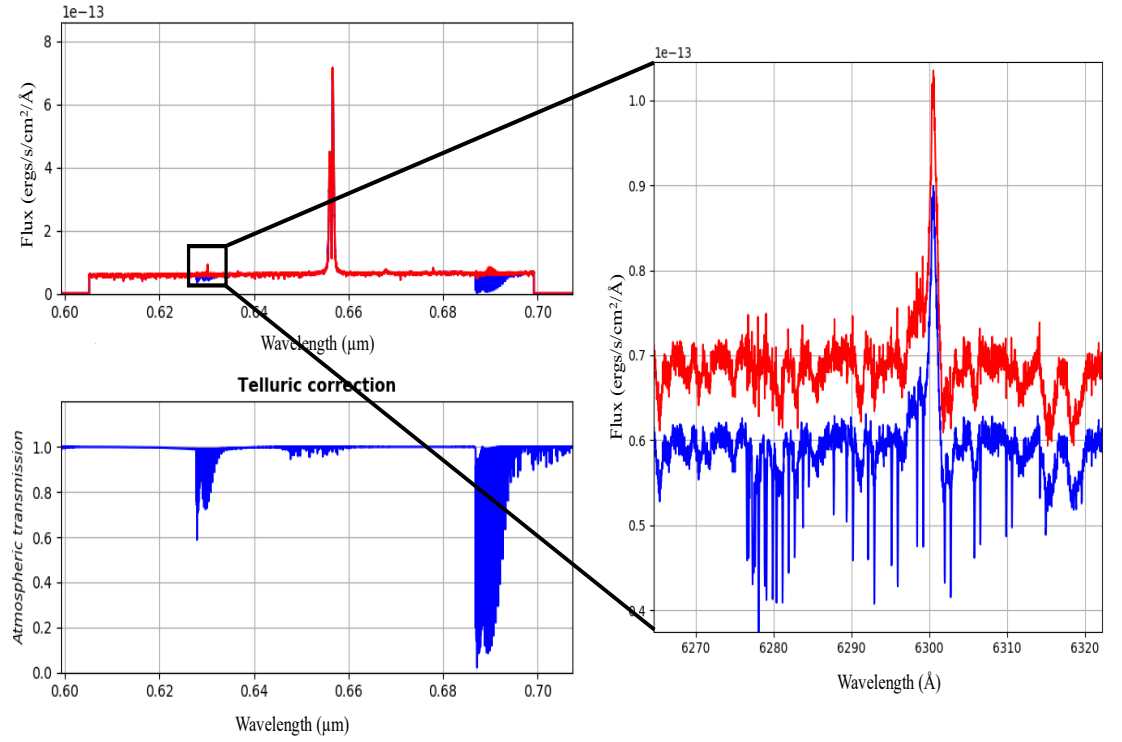


FIGURE 2.13: MOLECFIT telluric correction example on the [O I] $\lambda 6300$ of the CTTS CW Tau. Top left: The full spectrum from 6000 to 7000 . In blue is the input 1D science spectrum **before** telluric correction and in red is the science spectrum **after** correction. Bottom left: Shows the modelled atmospheric transmission spectrum. 1 = full transmission and 0 = full absorption. Left panel: Shows an zoom in of the [O I] $\lambda 6300$ line. This is one of the key lines typically investigated with SA that is most effected by telluric contamination. The corrected spectrum (in red) has been shifted up the y-axis to aid visualisation.

Next, the photospheric absorption features have to be removed from the science spectrum. A quantitative approach to correct this unwanted absorption was first introduced by Hartigan et al. (1989). In essence, this method compares the science spectrum of the target CTTS (accreting) to a template spectrum of a Weak Line T Tauri star (WTTS, no accretion) of the same spectral type. By this comparison, an estimate of the veiling imposed on the spectrum of the CTTS can be made by examining how the absorption features are "filled in" by accretion. This is summarised in Equation 2.6.

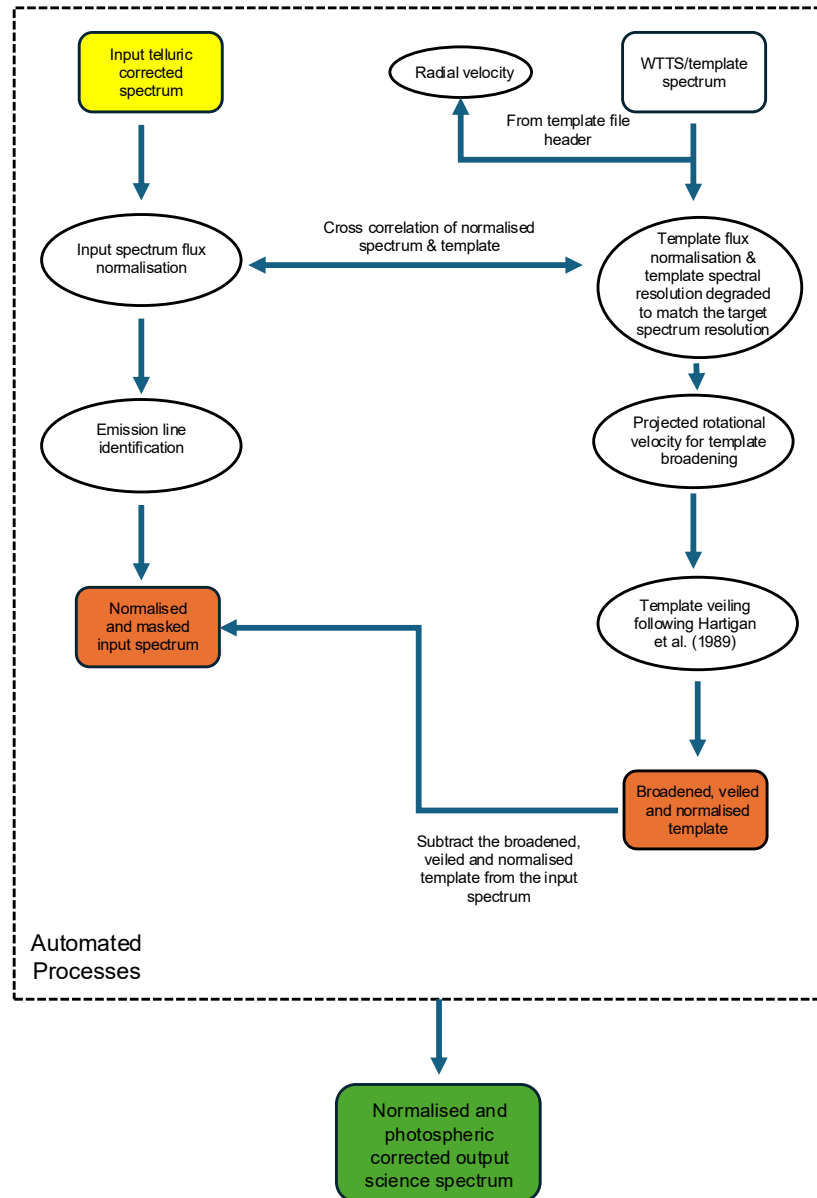
$$O(\lambda_i) = A[S(\lambda_i) + V(\lambda_i)] \quad (2.6)$$

where $O(\lambda_i)$ is the observed spectrum of the CTTS in wavelength interval λ_i , $S(\lambda_i)$ is the WTTS spectrum where only the photosphere is present, $V(\lambda_i)$ is the veiling spectrum (due to the excess in accretion between the CTTS spectrum and the WTTS spectrum) and A is a normalisation constant. The veiling quantity is calculated as several values over small wavelength intervals as it is difficult to measure veiling across the full range of the spectrum (Hartigan et al., 1989). These veiling quantities are commonly written as r_j and can be expressed by Equation 2.7. The veiling quantities can then be applied to the WTTS spectrum to scale its photospheric absorption features of the photospheric features present in the CTTS spectrum.

$$r_j = \frac{F_{excess}}{F_{photosphere}} \quad (2.7)$$

The WTTS spectrum is then rotationally broadened and shifted in velocity to match the width and position of the photospheric absorption features. The veiled, broadened and shifted WTTS template is then continuum normalised and subtracted from the normalised CTTS science spectrum to remove the photospheric absorption features. A WTTS of the same or similar spectral type must be used for this process as the exact shape of the photospheric features is dependent on temperature and stellar age (Stock et al., 2022). A *Python* based photospheric removal tool (PRT) has been developed within the STAR MELT *Python* package (Campbell-White et al., 2021) based on the method described in Hartigan et al. (1989). This PRT has been adopted for use

in this thesis. A schematic diagram of the STAR MELT PRT workflow is shown in Fig. 2.14.



Legend

Yellow = Input
 Orange = Intermediate Products
 Green = Output

FIGURE 2.14: The STAR-MELT PRT workflow. Steps within the dashed box are generally automated, but some user interaction may be required to obtain the best possible output.

The PRT comes pre-packaged with over 60 high spectral resolution WTTS template spectra with the option for the user to use their own template. Figure 2.15 shows an example of the templates of spectral type K7 available within the STAR MELT PRT.

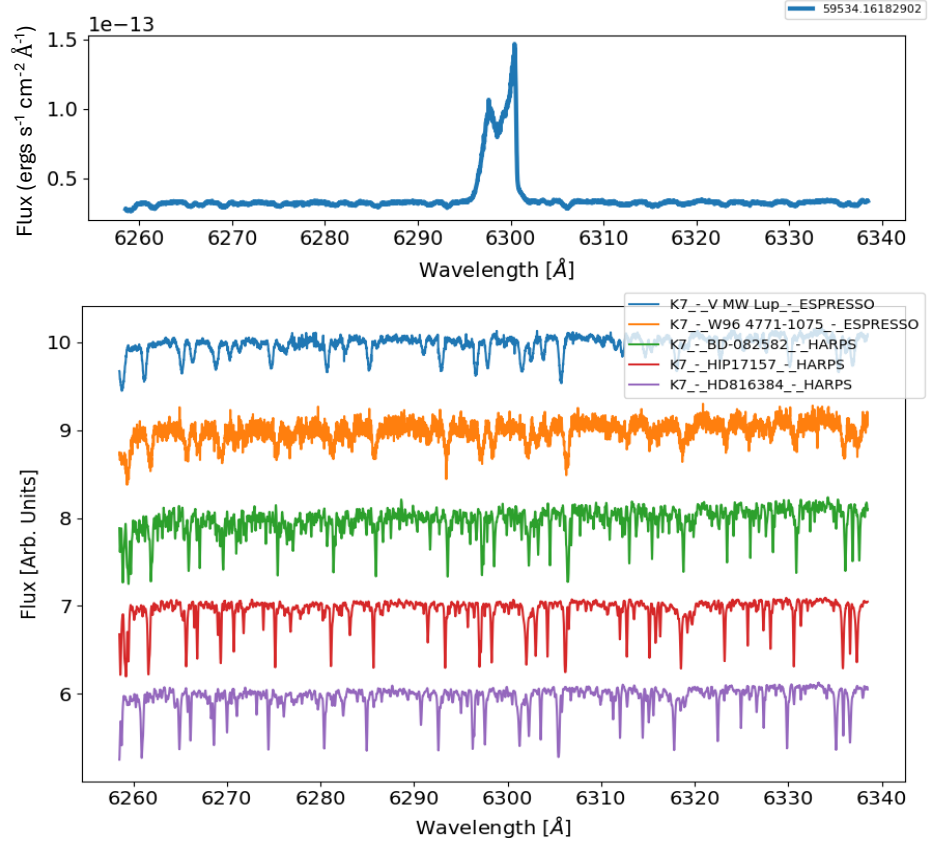


FIGURE 2.15: Example of STAR MELT input. Top: Input telluric corrected [O I] $\lambda 6300$ line of the CTTS DG Tau with a spectral type of K7. Bottom: The WTTS template spectra of type K7 available in STAR MELT to fit the photospheric absorption features.

Following the procedure described above, the PRT scales and aligns the photospheric features in the WTTS to match the photospheric features as closely as possible and then subtracts them. An example of a corrected and normalised [O I] $\lambda 6300$ line profile is presented in Fig. 2.16.

Heliocentric and radial velocity corrections

Once the telluric and photospheric corrections have been performed to the 1D spectra, each spectrum has to be shifted from the topocentric (telescope) frame of reference to

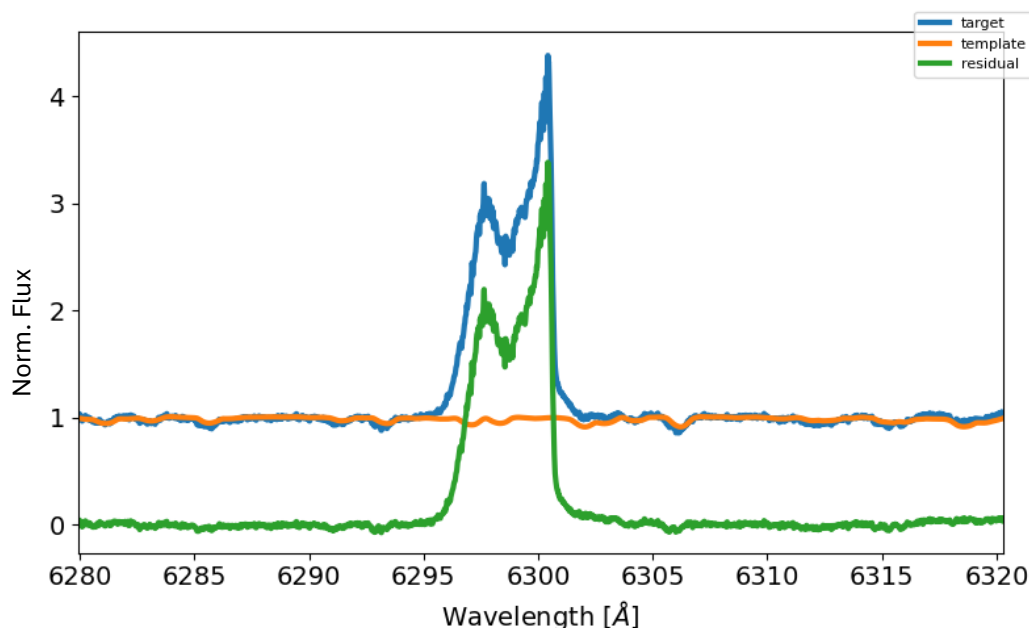


FIGURE 2.16: Example of STAR MELT output, showing the normalised [O I] $\lambda 6300$ line of DG Tau before (blue) and after (green) photospheric correction. The orange line is the fit of the veiled and broadened WTTS spectrum to the uncorrected line profile.

the stellocentric (stellar) frame of reference so that one can compare kinematic information from the spectra of different target stars OR multiple epochs of spectra of the same star.

In order to convert the line profiles from the topocentric frame into the stellocentric frame two velocity corrections were applied to each line profile. A heliocentric velocity correction and a stellar radial velocity correction. The heliocentric velocity correction converts from the topocentric frame of reference into the heliocentric (Sun) frame. This correction was calculated using the *radial_velocity_correction* function available within *Astropy* package in *Python* (Astropy Collaboration et al., 2013, 2018, 2022). Alongside this, the *rvcorrect* method available in the Image Reduction and Analysis Facility (IRAF) was also used to confirm the accuracy of the heliocentric velocity correction calculated using *Astropy*. Both functions take in information about the date and time of observation and the position of the target and telescope to calculate the correction factors to shift spectra into the heliocentric frame.

Secondly, a stellar radial velocity correction was calculated following Pascucci et al., 2015. I (the author of this thesis) have adapted this method into *Python* script. The method follows a Monte Carlo approach in which normally distributed random noise is added to the Li $\lambda 6707$ or Ca $\lambda 6439$ absorption line. After which the line is fit with a single inverted Gaussian function in the heliocentric reference frame and the peak velocity of the fit to the absorption line is recorded. The noise is then removed, and new normally distributed noise is added to the line. The line is refit to obtain another value of peak velocity. This process is repeated one thousand times in total. This was done to evaluate the uncertainty in the peak velocity of these absorption features. If no noise was added to the Li $\lambda 6707$ or Ca $\lambda 6439$ line profiles, the uncertainty on the fit to each line profile alone would likely provide an underestimate of the uncertainty on the peak velocity as the fit uncertainty estimate alone doesn't account for the observing conditions, spectral resolution or the signal-to-noise of the data. One thousand iterations was chosen as to obtain a statistically significant sample of fits to each absorption feature from which, an accurate standard deviation and hence, an uncertainty on the radial velocity could be obtained. The heliocentric stellar radial velocity is then the average value of the one thousand peak velocities that were obtained from this fitting procedure. An example is shown in Fig. 2.17.

The uncertainty on the heliocentric radial velocity measurement is the standard deviation (σ) of the one thousand peak velocities plus the value, σ_{fit} , derived by Equation 2.8. Where S/N in this case represents the signal-to-noise ratio of the absorption line.

$$\sigma_{fit} = \frac{0.4 \times seeing \text{ (in arcseconds)}}{S/N} \quad (2.8)$$

The radial velocity of each star to be presented in Chapter 4 is shown in Table 2.2.

2.2.3 Post pipeline data processing: 2D spectra

The 2D spectra also require post processing to make them suitable for analysis. This is particularly important with SA in mind as any imperfection in the 2D spectra may

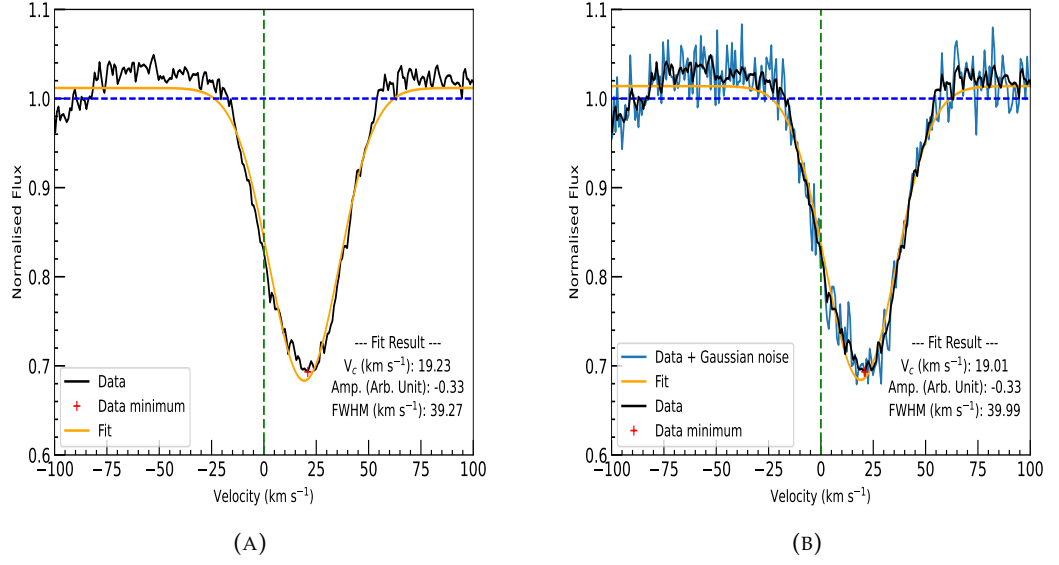


FIGURE 2.17: An example of calculating heliocentric radial velocity via the method of Pascucci et al. (2015). (A): Fit to the Li $\lambda 6707$ line of DG Tau before adding noise. (B): Fit after adding noise. The green and blue dashed lines correspond to 0 km s^{-1} and the normalised continuum level respectively.

TABLE 2.2: Radial Velocity values for each target star of Chapter 4 calculated using the method of Pascucci et al. (2015). The Li $\lambda 6707$ and Ca $\lambda 6439$ lines have been used to calculate a radial velocity value for each target star. These values are compared to each other and to estimates in the literature.

Source	RV_{Ca} [km s^{-1}]	RV_{Li} [km s^{-1}]	RV_{lit} [km s^{-1}]	Ref.
TW Hya	12.4 ± 0.3	12.1 ± 0.3	13.55 ± 0.34	F18
T Cha	16.3 ± 0.6	17.3 ± 0.5	17.3	GDR3
MP Mus	12.3 ± 0.4	12.5 ± 0.4	10.7	P20
LkH α 330	16.7 ± 0.7	15.4 ± 0.9	19.0 ± 0.5	B19
GQ Lup	-3.3 ± 0.3	-1.7 ± 0.3	-2.13 ± 0.29	F18
DG Tau	18.3 ± 0.7	19.0 ± 0.5	14.6 ± 0.5	B19
CW Tau	12.3 ± 0.6	9.4 ± 0.4	15.0 ± 0.5	B19
GK Tau	13.2 ± 0.4	14.8 ± 0.3	18.52 ± 0.64	F18
DK Tau	16.2 ± 0.4	17.7 ± 0.4	16.60 ± 0.42	F18
EX Lup	1.8 ± 0.5	1.4 ± 0.4	2.0 ± 1.0	F18

B19 = Banzatti et al. (2019); F18 = Fang et al. (2018); GDR3 = Gaia Data Release 3 (Prusti et al., 2016; Vallenari et al., 2023); P20 = Pascucci et al. (2020)

produce a false signal when performing SA. Two post processing steps are performed on the 2D spectra once they are initially reduced by the UVES pipeline. These are one,

skyline (and cosmic ray pixel) removal and two, continuum subtraction. These steps are now described below.

Skyline and cosmic ray removal

Skylines are the 2D equivalent of telluric features. They appear as narrow vertical stripes across the whole spatial range of the 2D spectrum as the telluric features come from the whole sky (see Fig. 2.19, top panel). Cosmic rays are high energy particles that strike the CCD causing pixels in the detector to saturate. Unfortunately, the current version of MOLECFIT is unable to handle 2D data so, both skylines and cosmic ray effected pixels were removed using the *splot* task in IRAF along each row of the dispersion axis in a line-by-line fashion.

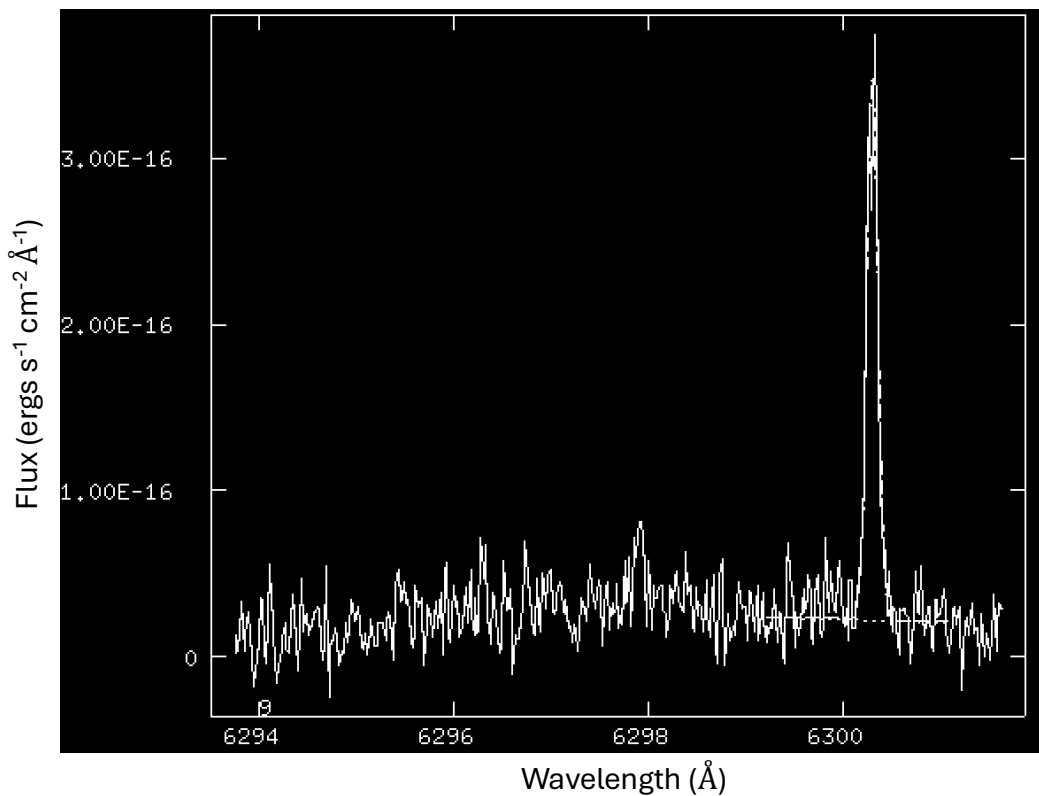


FIGURE 2.18: Example of fit to skyline from one row along the dispersion axis of a 2D spectrum. This row is a cut along the 2D spectrum in a row that is far from the centre of the continuum. The spectrum cut is the solid white line. The Gaussian fit is denoted by the dotted white line.

Figure 2.18 presents an example of one row along the dispersion axis. The *splot* task in

IRAF was used to fit a Gaussian function to a skyline effected region in this row along the dispersion axis of the 2D spectrum. This is done in all rows along the dispersion axis until the skylines and cosmic ray effected pixels are removed. Figure 2.19 shows an example of a 2D spectrum before and after skyline removal. A similar procedure is used to remove cosmic ray effected pixels in the 2D spectrum.

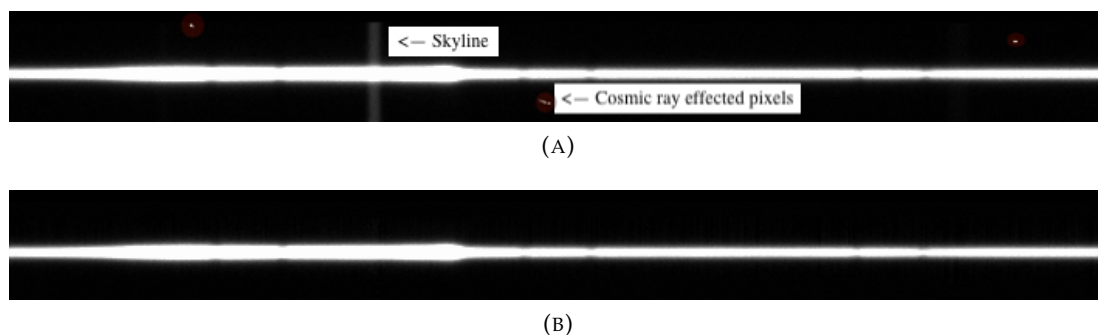


FIGURE 2.19: The 2D spectrum of the [O I] $\lambda 6300$ line of DG Tau. (A): before and (B): after cosmic ray and skyline removal. The cosmic ray effected pixels have been marked with semi transparent red ellipses.

Continuum subtraction

Continuum subtraction was performed on the 2D spectra for the reasons outlined earlier in this chapter. A custom *Python* script was provided for this thesis by Dr. A. Murphy (private communication) to perform the continuum subtraction on each 2D spectrum. This script takes skyline corrected 2D spectrum from the previous step as input. First, the central position of the continuum is estimated via a Gaussian fit to cross section cut, perpendicular to the dispersion axis (see Fig. 2.20).

Once the continuum centre is found, the script extracts a sample 1D spectrum by median combining a few rows either side of the continuum centre. The exact number of rows either side of continuum centre to be combined is set via the user. An example of such an extracted spectrum is demonstrated in Fig 2.21.

Using a figure akin to Fig. 2.21, regions in pixel space, free from line emission or absorption features were selected as continuum regions to be used to estimate the continuum level in the spectrum. Some key ELRs (i.e. the ELRs to be studied after

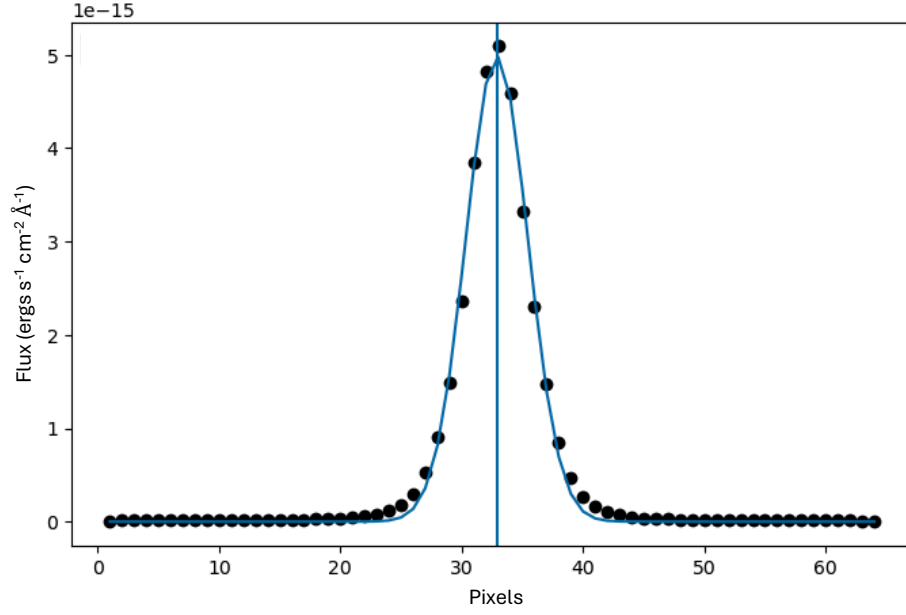


FIGURE 2.20: Gaussian fit to the profile extracted perpendicular to the dispersion axis to find where the continuum is centred. The point at which the blue line intersects the x-axis corresponds to the estimate of the position of the continuum centre in pixel space.

data processing) were also selected to ensure they were masked and excluded from the continuum removal procedure. The continuum is then estimated from these selected continuum regions by fitting them with a polynomial of degree three. The resulting polynomial fit is then subtracted from every row that was selected by the user (i.e. the rows that were selected to make the median combined spectrum in Fig. 2.21) to remove the continuum emission from the 2D spectrum. An example of the output from this continuum subtraction routine is presented in Fig. 2.22).

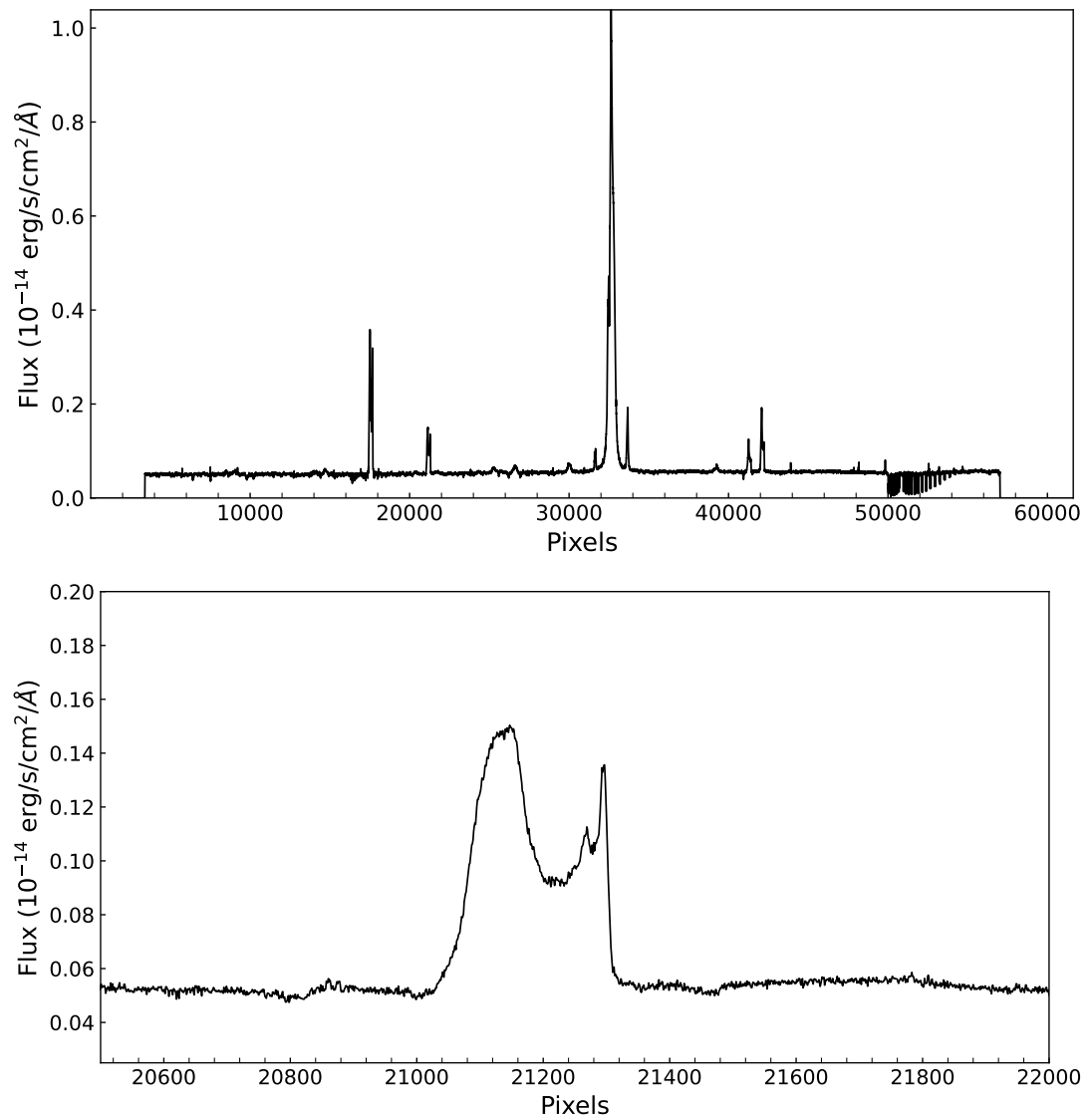


FIGURE 2.21: Example of extracted spectrum from continuum subtraction script. Top: Full length extracted spectrum in pixel space. Bottom: A zoomed region into a user specified region in pixel space.

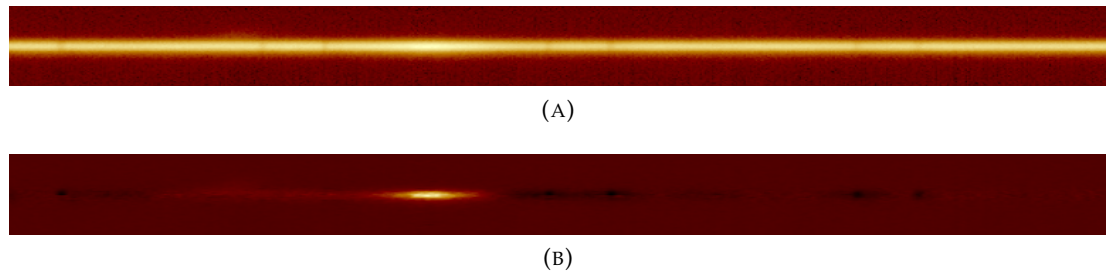


FIGURE 2.22: The 2D spectrum of the [O I] $\lambda 6300$ line of DG Tau. (A): before and (B): after continuum subtraction.

Chapter 3

Interferometry with the Atacama Large mm/sub-mm Array

3.1 Beyond the optical

As outlined in Chapter 1, very young and embedded YSOs are not optically visible due to extinction caused by the material of the surrounding natal envelope. As such, observations at longer wavelengths (Fig. 3.1) are required to observe and understand the formation processes of the youngest and most embedded YSOs and young BDs.

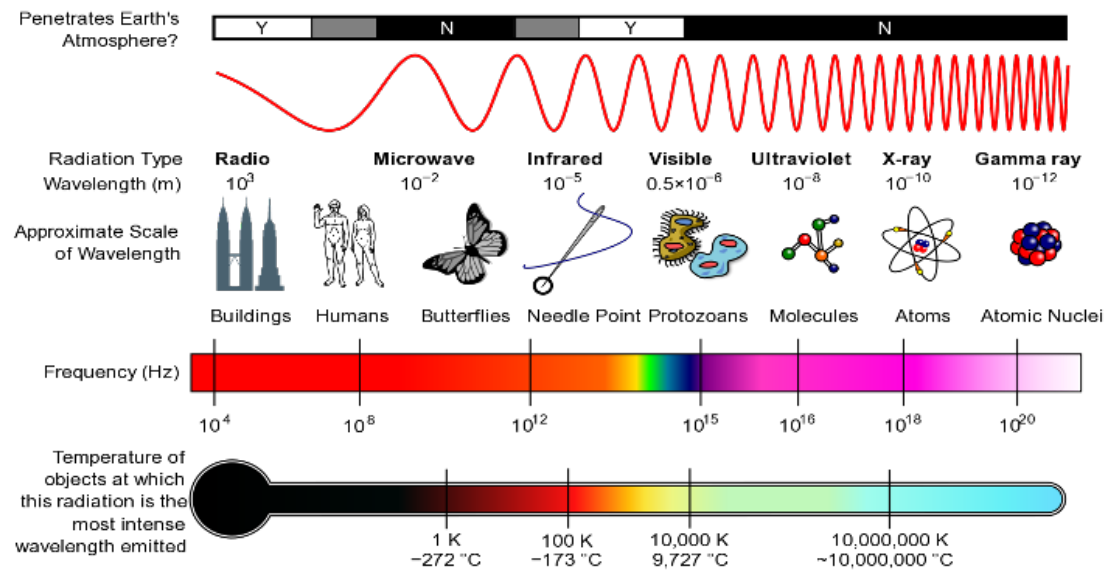


FIGURE 3.1: The electromagnetic spectrum. Reproduced under the Creative Commons Attribution-Share Alike 3.0 license Wikipedia contributors (2025).

This chapter focuses on a sub-branch of radio astronomy called millimetre/sub millimetre (mm/sub-mm) astronomy. The sub-mm branch is considered the highest frequency (and therefore shortest wavelength) of light in the radio spectrum. It spans

from around 35 gigahertz (GHz, 8.6 mm) to about 1 terahertz (THz, 0.3 mm). Observations at these wavelengths have provided a wealth of information in regard to both outflow and accretion processes in YSOs and young BDs (Pineda et al., 2020; Santamaría-Miranda et al., 2021; Yen et al., 2019). The rest of this chapter is divided as follows. First, it provides an overview on radio astronomy, the ALMA telescope and the interferometric method used to obtain the results presented in Chapter 6. Lastly, this chapter concludes with a section on working with the output science data of ALMA.

3.2 Brief history of radio astronomy

The origins of radio astronomy can be traced back to the discovery of radio emission from the Milky Way by Karl Jansky in 1933 (Jansky, 1933). Following this was the work of Grote Reber who used his own 30-foot (≈ 9 m) antenna to map radio emission from the galaxy at 160 megahertz (MHz). He identified several secondary peaks in radio emission in addition to the strong radio emission from the centre of the Milky Way. One was the supernova remnant Cas A and another was the radio galaxy Cyn A (Reber, 1940, 1944). Following the second world war, a number of developments and discoveries were made. In 1946, Martin Ryle and D.D. Vonberg made the first step in the development of high resolution radio observations by making the first interferometric observation with a pair of radio antennas (Ryle and Vonberg, 1946). In the 1950s, the now famous 21 cm line was used to map the distribution of hydrogen gas throughout the Milky Way (Ewen and Purcell, 1951). Alongside this, a deeper theoretical understanding was reached on how synchrotron radiation is produced via the emission of photons from relativistic electrons as they spiral around magnetic fields (Alfvén, 1954). During the 1960s, radio emission led to the discovery of pulsars and the cosmic microwave background, both of which were awarded the Nobel Prize in physics in 1978 and 1974. The modern era of radio imaging began in the late 60s into the mid 70s with the first very long baseline interferometer (Ryle, 1975) and in 1974 with the development of the aperture synthesis technique and the CLEAN algorithm by Högbom (1974) and Hogbom and Brouw (1974). These techniques have

been refined over the last five decades to unearth a myriad of information on different astrophysical processes. From molecular outflows launched from YSOs (Santamaría-Miranda et al., 2020), the imaging of sub-structures in circumstellar discs (Andrews et al., 2018; Hughes et al., 2025), applications in astrochemistry (Artur de la Villarmois, 2023), water maser emission (de Gregorio-Monsalvo et al., 2023) to the first image of a black hole M87 by the Event Horizon Telescope (EHT) (Event Horizon Telescope Collaboration et al., 2019). Given all the above, it is clear that radio astronomy is a vibrant and active field in modern astronomy.

3.3 Radio interferometry basics

The wavelength (λ) of radio signals are hundreds to millions of times longer than optical wavelengths and as such, the angular resolution is severely diffraction limited for a single radio antenna of diameter D and crude by optical standards (Equation 3.1). The technique of interferometry was developed and used to improve the angular resolution of observations at radio wavelengths. The fundamental idea behind interferometry is that there exists a Fourier transform relationship between the radio sky brightness distribution and the response of the interferometer. By measuring the so-called complex visibilities in Fourier space, a reconstructed image of the radio sky brightness distribution can be made by means of a Fourier transform (FT) on the complex visibilities (Stanimirovic, 2002). Interferometry involves the combination of signals from two or more physically separated antennas. The signals received from each antenna are interfered with each other. This then allows a given radio sky brightness distribution to be sampled at an angular scale much smaller than the crude diffraction limit provided by a single antenna.

$$\theta_{res} \propto \frac{\lambda}{D} \quad (3.1)$$

Prior to the discussion on interferometry, it is pertinent to first discuss the response of a single antenna to incident radiation.

3.3.1 Single antenna response

Fig. 3.2 shows a normalised response of an antenna to incident radiation as a function of the angles θ and ϕ . The solid angle of the antenna beam in steradians is the beam power pattern integrated over the whole sky (Equation 3.2). The solid angle of the antenna beam integrated over the main beam only is given in Equation 3.3.

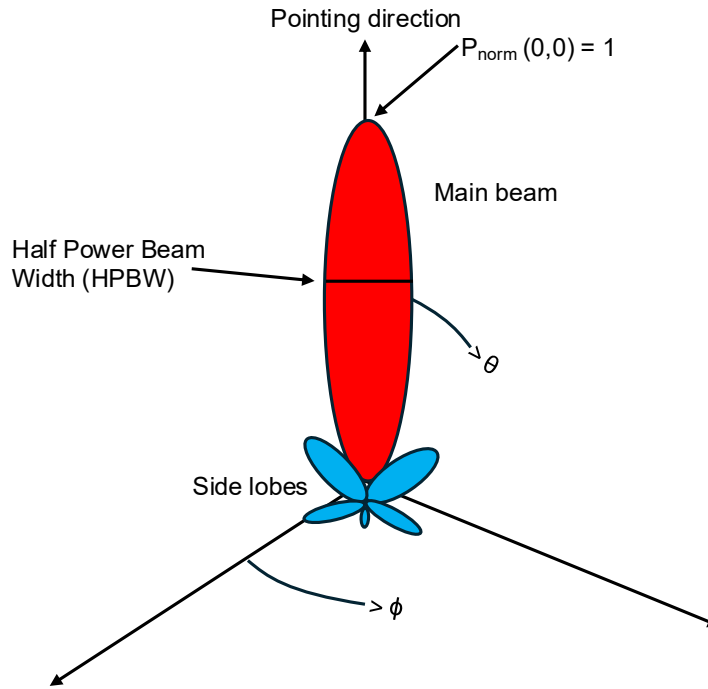


FIGURE 3.2: Example of a normalised beam power pattern.

$$\Omega_a = \iint_{4\pi} P_{norm}(\theta, \phi) d\Omega \quad (3.2)$$

$$\Omega_{mb} = \iint_{main\ beam} P_{norm}(\theta, \phi) d\Omega \quad (3.3)$$

In an idealised situation for a 100% efficient antenna, the effective aperture area of the antenna is given according to the antenna theorem (Stanimirovic, 2002), shown by Equation 3.4.

$$A_e = \frac{\lambda^2}{\Omega_a} \quad (3.4)$$

In reality, most of the power received by the antenna enters through the main beam and only a fraction of the incident power on the antenna is received through the beam side lobes. The efficiency of the main beam to incident power is then defined as the ratio of the main beam solid angle to the solid angle subtended by the whole antenna. Correspondingly, the efficiency of the antenna aperture to incident power is therefore the ratio of the effective aperture area (A_e) to the geometric area of the antenna.

At radio wavelengths, the source of radiation is treated like a Plank blackbody that radiates as a temperature T . As such, the spectral radiance of a blackbody at a particular frequency ν is given by the Plank blackbody function. At these long wavelengths, the Rayleigh-Jeans approximation holds (i.e. $h\nu \ll kT$). The brightness emitted by the blackbody source can therefore be approximated as proportional to its thermodynamic temperature. So, the Plank blackbody function can be written in the simplified form shown in Equation 3.5, where k is the Boltzmann constant, h is the Plank constant and c is the speed of light.

$$I_\nu(T) = \frac{2\nu^2}{c^2}kT = \frac{2}{\lambda^2}kT \quad (3.5)$$

Rearrangement of the above equation then gives an expression for the temperature of a blackbody that emits at the same intensity as the source observed by the antenna (Equation 3.6). This is referred to as the brightness temperature T_b .

$$T_b = \frac{\lambda^2}{2k}I_\nu(T) \quad (3.6)$$

In radio astronomy the antenna is treated as a resistor that heats up in response to the incident power. According to the Nyquist theorem (Nyquist, 1928), if that antenna receives a certain power P_ν we can define an antenna temperature T_A that yields the same power (see Equation 3.7)

$$P_\nu = kT_A \quad (3.7)$$

In Fig. 3.3, a radio source subtends a solid angle of Ω_s on the sky. In the simplified

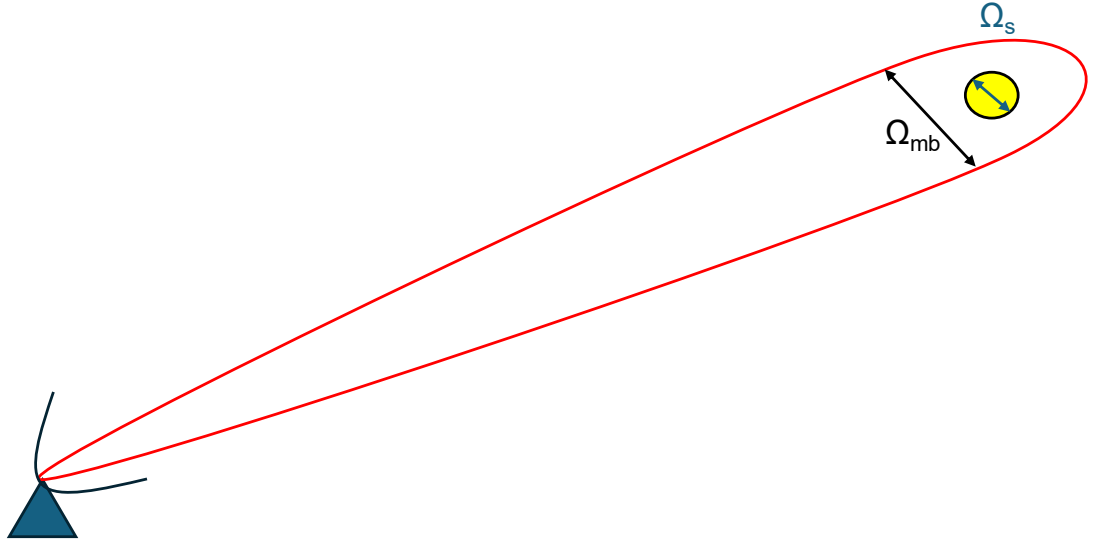


FIGURE 3.3: Schematic diagram of single antenna main beam and point-like source. The main beam of the antenna and the source being observed subtend angles of Ω_{mb} and Ω_s on the sky respectively.

case where the source is of uniform brightness (i.e. constant brightness temperature), the flux density S_ν is given by Equation 3.8.

$$S_\nu(T) = \iint_{source} I_\nu(T) d\Omega = \frac{2k}{\lambda^2} T_{const} \Omega_s \quad (3.8)$$

The power per unit bandwidth (i.e the flux density of the source times the effective aperture area) received by an antenna with effective aperture area A_e and a beam pattern $P_{norm}(\theta, \phi)$ from a source with a brightness distribution $I_\nu(\theta, \phi)$ on the sky is given by Equation 3.9.

$$P_\nu = \frac{1}{2} A_e \iint_{source} I_\nu(\theta, \phi) P_{norm}(\theta, \phi) d\Omega = kT_A \quad (3.9)$$

The factor one half comes from the fact that the antenna receiver is only sensitive to one type of polarization. ALMA is designed to be sensitive to both linear and circular polarization, so this factor is 1. Using Equation 3.6, the above equation can be rewritten in terms of the source brightness temperature distribution $T_b(\theta, \phi)$:

$$P_\nu = \frac{A_e k}{\lambda^2} \iint_{source} T_b(\theta, \phi) P_{norm}(\theta, \phi) d\Omega \quad (3.10)$$

Remember that the received power is proportional to the antenna temperature ($P_v = kT_A$), so the antenna temperature T_A can be written in terms of the brightness temperature of the source and the antenna beam power pattern.

$$T_A = \frac{A_e}{\lambda^2} \iint_{source} T_b(\theta, \phi) P_{norm}(\theta, \phi) \quad (3.11)$$

The increase in temperature of the antenna due to the incident radiation received by the antenna from the observed source is proportional to the flux density of the source (i.e. $P_{rec.} = A_e S_v$). Hence, the observed flux density of the source by the antenna is given by:

$$S_{obs} = \iint_{source} I_v(\theta, \phi) P_{norm}(\theta, \phi) d\Omega \quad (3.12)$$

Equation 3.12 shows that observed flux density is the brightness distribution of the observed source convolved with the power pattern of the antenna beam (i.e. the antenna response). There is a similar relationship between the sky brightness distribution and the response of an interferometer which consists of more than one antenna. This is discussed in the next subsection.

3.3.2 The two-element interferometer and aperture synthesis

As stated previously, interferometry combines signals from two or more physically separated antennas to obtain angular resolution better than the diffraction limit of a single antenna. This technique is sometimes called aperture synthesis. In essence, it is the synthesising of a large pseudo-antenna from the combination of signals from an array of smaller antennas.

In simplest case of a two antenna (or two element) interferometer separated by a baseline vector $\mathbf{b}$. The length of the baseline can be measured in units of observing wavelengths λ . The length of the the baseline vector $\mathbf{b}$ can also be written as $b = L/\lambda$ where L is the distance between the two antennas. Figure 3.4 shows a schematic of a two element interferometer with such a setup. Both antennas observe a position on the sky s_0 located at an angle θ from the meridian. The projected separation of the two antennas

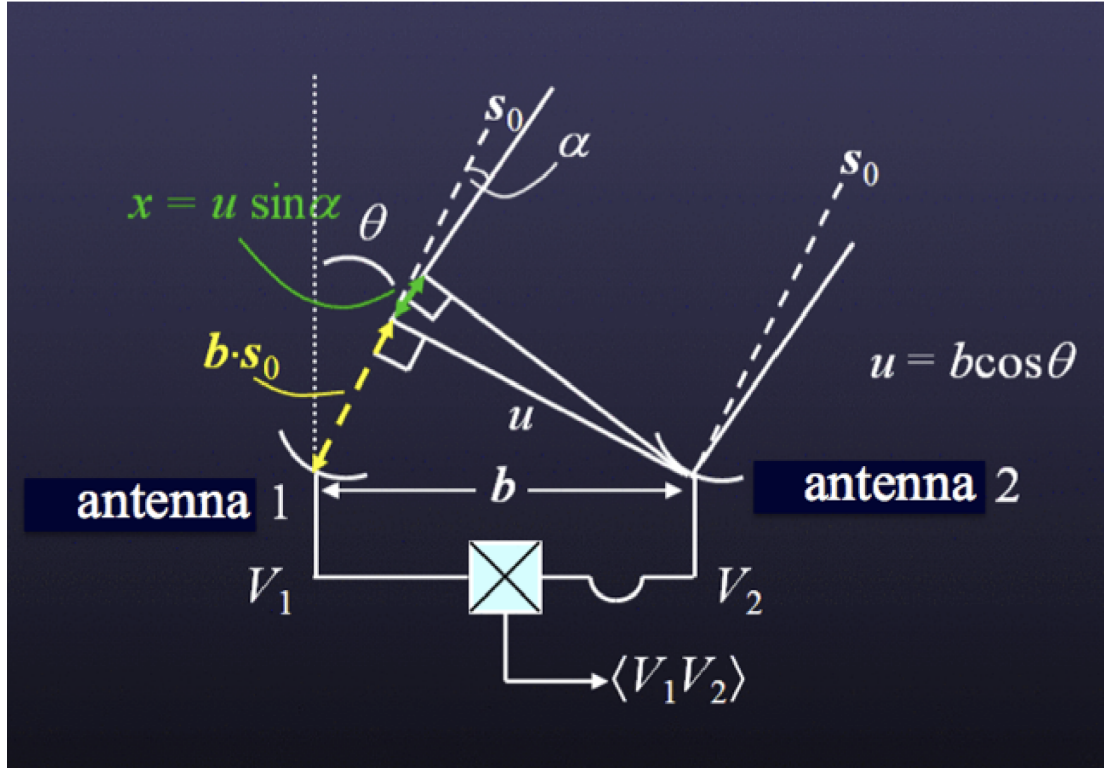


FIGURE 3.4: Taken from the ALMA Cycle 12 technical handbook. Available at: <https://almascience.eso.org/documents-and-tools/cycle12/alma-technical-handbook>. An ideal 1D interferometer consisting of two antennas, 1 and 2, separated by physical distance (i.e., a baseline) b . The antennas are both pointed towards a sky location given by s_0 , which is at an angle θ from the meridian. The projected distance between the two antennas in that direction is thus $u = b \cos(\theta)$. The two antennas are connected to a correlator where the voltages detected from each antenna are combined.

from the perspective of the observed source is $u = b \cos(\theta)$. An on axis electromagnetic wavefront from the source will reach antenna 2 before antenna 1 since the wavefront has to travel an extra distance (or path length) of $b \cdot s_0 = b \sin(\theta)$ to reach antenna 1. This means that antenna 1 experiences a geometrical time delay relative to the time the wavefront is received by antenna 2. This time delay is $\tau_g = b \cdot s_0 / c$ and can be compensated for (e.g. electronically) by inserting a delay in the signal path of antenna 2 such that the signal from both antennas arrive at the correlator with the same phase.

In the case where the incoming wavefront is slightly off axis (by a small angle α), the 1D sky position of the off axis wavefront can be described as an angle $l = \sin(\alpha)$ (also known as a direction cosine). The off axis wavefront travelling to antenna 1 will

have to travel an extra path length to reach there compared to the off axis wavefront travelling to antenna 2. This extra path length is $x = u \sin(\alpha) = ul$. The distance x can be written in units of λ such that x is a number of wavelengths in a given distance. This extra path length results in phase differences (or delays) between the two antennas. As such, the voltage response of antenna 2 (V_2) can be written as a product of the voltage response of antenna 1 (V_1) and a phase delay factor that varies sinusoidally as a function of the off axis angle α (see Equation 3.13).

$$V_2 = V_1 e^{2\pi i(ul)} \quad (3.13)$$

Expanding this argument to two dimensions introduces another off axis direction on the sky called β that is orthogonal to α . The baseline vector $\mathbf{b}$ becomes 2D with components b_1 and b_2 in each direction. Also, $m = \sin(\beta)$ is this analogue to l in this new axis. The projected separation of the two antennas in 2D relative to the perspective of the observed source is then defined as $u = b_1 \cos(\theta)$ (as in the 1D case) and $v = b_2 \cos(\phi)$, where ϕ is the angle of the source position s_0 relative to the reference position. ϕ is the orthogonal equivalent to θ in this added dimension. Lastly, the extra path length for the off axis wavefront as it travels in this new direction is $y = v \sin(\beta) = vm$ (This is the equivalent of x in this added dimension). With these changes the voltage response of antenna 2 changes to (i.e. Equation 3.13 changes to):

$$V_2 = V_1 e^{2\pi i(ul+vm)} \quad (3.14)$$

where l and m are the direction cosines of each axis relative to a reference position in the East-West and North-South directions respectively (i.e. the image plane). For an on axis position s_0 has $l = 0$ and $m = 0$ and this is called the *phase centre*. u and v are called the *spatial frequency* components of the sinusoidal interference pattern (see Fig. 3.5 for an example) in the East-West and North-South directions respectively. These values are the projected lengths of the antenna separations measured in units of the wavelength at the time of observation.

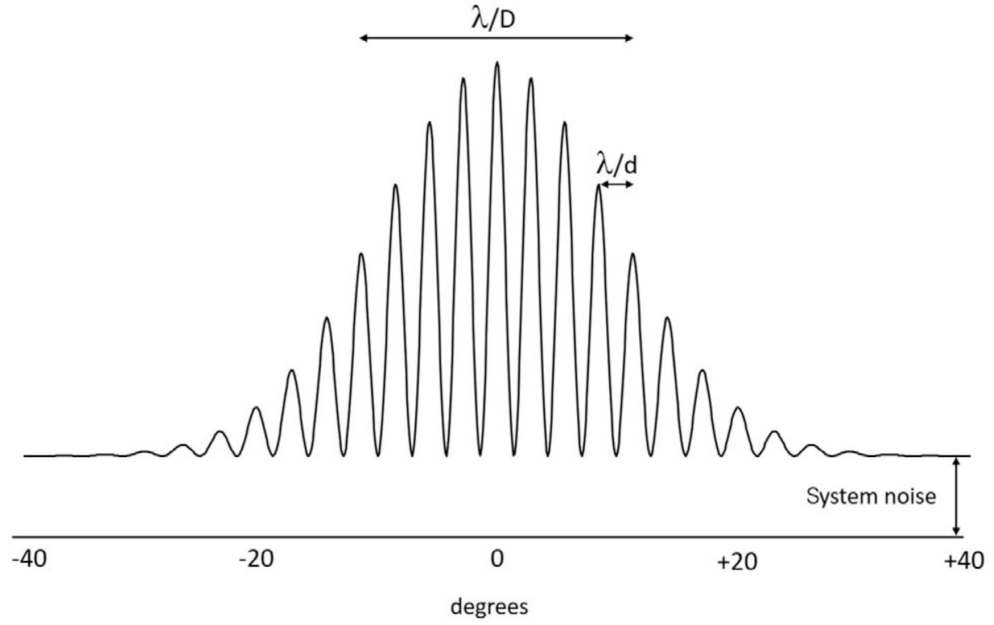


FIGURE 3.5: Taken from Burke, Graham, and Wilkinson (2019). Example polar diagram of a sinusoidal interference pattern from an "adding" interferometer consisting of two ideal and identical antennas. D = diameter of each identical antenna separated by a spacing of d (d here is the same as L in the notation above). In this example the antenna diameters were 2.4λ and the interferometer spacing was 20λ .

The signal correlator in this example acts as a multiplying and time averaging device for the signals coming from antennas 1 and 2. With the assumption that signals emanating from different parts of the sky are incoherent (i.e., their phases are uncorrelated), the time averages of the correlation of those signals will be zero. Also, the power received by the antenna is also proportional to I_ν (see Equation 3.9). Hence, Equation 3.15 defines the output of the signal correlator as follows:

$$\langle V_1 V_2 \rangle \propto \iint I(l, m) e^{2\pi i(ul + vm)} dl dm \quad (3.15)$$

where $I(l, m)$ is the brightness distribution on the sky. The correlator therefore measures a quantity known as a *complex visibility* $\mathcal{V}$. By definition, this quantity is the Fourier transform of the brightness distribution on the sky:

$$\mathcal{V}(u, v) = \iint I(l, m) e^{2\pi i(ul + vm)} dl dm = A e^{i\phi} \quad (3.16)$$

The complex visibility can be written as a complex exponential with an amplitude

term A , which describes the source brightness and a phase term φ , which describes the source position relative to the phase centre at spatial frequencies u and v .

In summary, the recovered image of the sky brightness distribution from the process of interferometry is the "sum" (i.e Fourier transform) of all the complex visibilities. Each complex visibility represents a point on the u - v plane and has an amplitude and phase representing the brightness and relative position of emission from the observed source on a specific angular scale. The angular scale that is sampled by a pair of antennas in the interferometer is dependent on the fringe spacing in the sinusoidal interference pattern (see Fig. 3.5) obtained when the signals are correlated together. The on-sky fringe spacing in the 1D case is given by:

$$\text{Fringe Spacing} = 1/u \text{ (in radians)} = 1/(b \cos(\theta)) = \lambda/(L \cos(\theta)) \quad (3.17)$$

Equation 3.17 then asserts that the angular scale of the on sky emission that is "seen" (or sampled) by the interferometer is proportional to the observing wavelength λ , and inversely proportional to the physical separation between a pair of antennas (see Equation 3.18). Therefore, to recover a true representation of the on sky emission, one many complex visibility values must be aquired to well sample the u - v (or Fourier) plane. A pair of antennas at a given baseline and observing wavelength generates two points on the u - v plane as these complex visibility values are complex value functions with Hermitian symmetry (i.e. they are inversion symmetric on the u - v plane), thereby obtaining the value of $\mathcal{V}(u, v)$ and the complex conjugate $\overline{\mathcal{V}}(u, v)$ for the same antenna setup. See Fig. 3.6 for an example.

$$\theta_{res} \propto \frac{\lambda}{b} \quad (3.18)$$

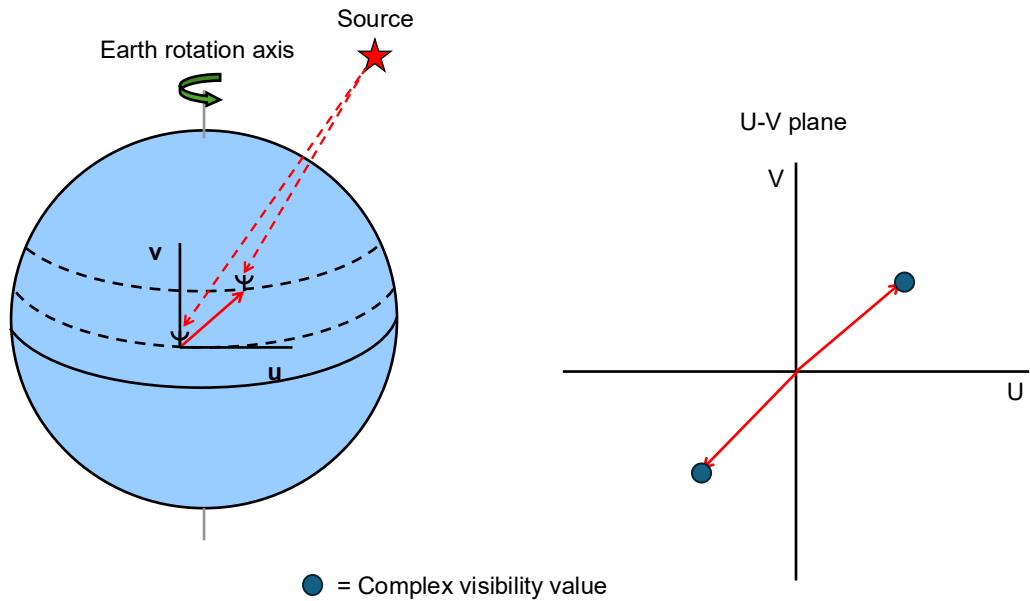


FIGURE 3.6: Schematic overview of how two antennas (one baseline) adds two points to the u-v (Fourier) plane. This figure is adapted with permission from I. de Gregorio Monsalvo.

In order to get an accurate reconstruction of the source brightness distribution the U-V plane must be well sampled to recover emission from as many angular scales as possible. There are several strategies to improve U-V plane sampling.

1. Add more antennas to the array. For N antennas, there is $N(N-1)/2$ independent baselines (i.e. not counting complex conjugates) and hence many more points sampling the u-v plane.
2. Earth's rotation helps sample the u-v plane. As the earth rotates, the projected distance to each pair of antennas changes from the perspective of the source. This changes the projected baseline length from the perspective of the source and hence, this traces out an arc of points on the u-v plane.
3. As hinted in Equation 3.17, the angular scale of the emission sampled by the interferometer is proportional to the observing wavelength λ . As such, observing at multiple wavelengths changes the fringe spacing in each of the interference patterns for each pair of antennas, so this effectively changes how the u-v plane is sampled. So, observing a target at multiple wavelengths is recommended.

Figure 3.7 presents an example of the u - v plane coverage for a 7m array ALMA observation taken with the *plotms* task in the Common Astronomy Software Applications (CASA) package.

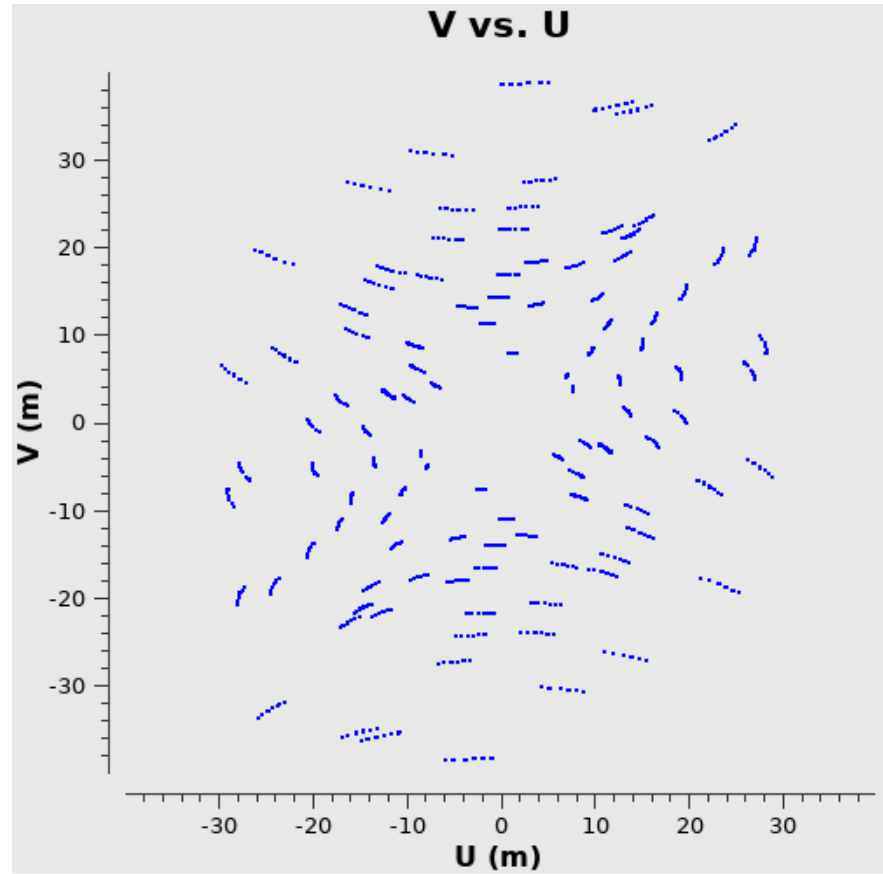


FIGURE 3.7: ALMA 7m array u - v plane coverage example. Also known as a sampling pattern.

3.4 An introduction to ALMA

ALMA is an array of 66 high precision antennas working together as an interferometer (Fig 3.8). The ALMA array consists of three sub arrays.

1. The 12m array: 50 12m diameter antennas used for high angular resolution observations.
2. The 7m array: 12 7m diameter antennas.
3. The Atacama Compact Array (ACA): An array of 4 12m antennas spaced closely together at the centre of the overall ALMA array.

The 7m array combined with the ACA is referred to as the Morita array in remembrance of Prof. Koh-ichiro Morita. The Morita array is used for the observation of extended sources.



FIGURE 3.8: The Atacama mm/sub-mm Array

ALMA is located at the Chajnantor plateau, in the Atacama desert of northern Chile ~ 5000 m above sea level. This site was chosen as mm/sub-mm observations are not unlike optical, in the sense that they are perturbed by the presence of water vapour and atmospheric turbulence. Selecting such a high and dry observing site minimises

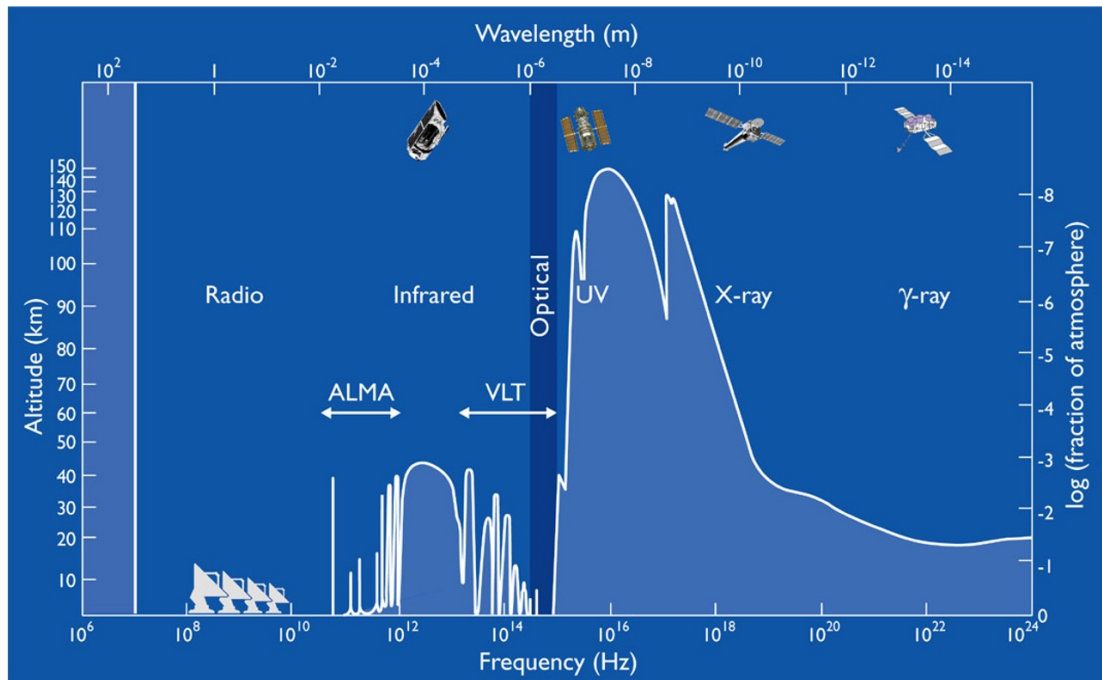


FIGURE 3.9: A typical atmospheric transmission curve at different wavelengths. Reproduced with permission from I. de Gregorio Monsalvo.

these effects. A typical atmospheric transmission curve is presented in Fig. 3.9. In this figure, the wavelengths or "bands" at which ALMA observes are selected to maximise the transmission of radiation through the atmosphere at these longer wavelengths.

More than twenty countries are involved in the ALMA project. These countries are divided into three executives (or ALMA Regional Centres ARCs), Europe, North America and East Asia, which individually act as interfaces to the ALMA user community. Each ARC provides user support, handles data delivery to each successful ALMA principle investigator and house a mirror of the ALMA data archive. Chile is the hosting country and as such, hosts the antennas and the Joint ALMA Observatory (JAO). The JAO is responsible for a multitude of operations, from scheduling and undertaking observations, to data quality assurance, data reduction, pipeline maintenance and delivery of science ready data to the ALMA data archive.

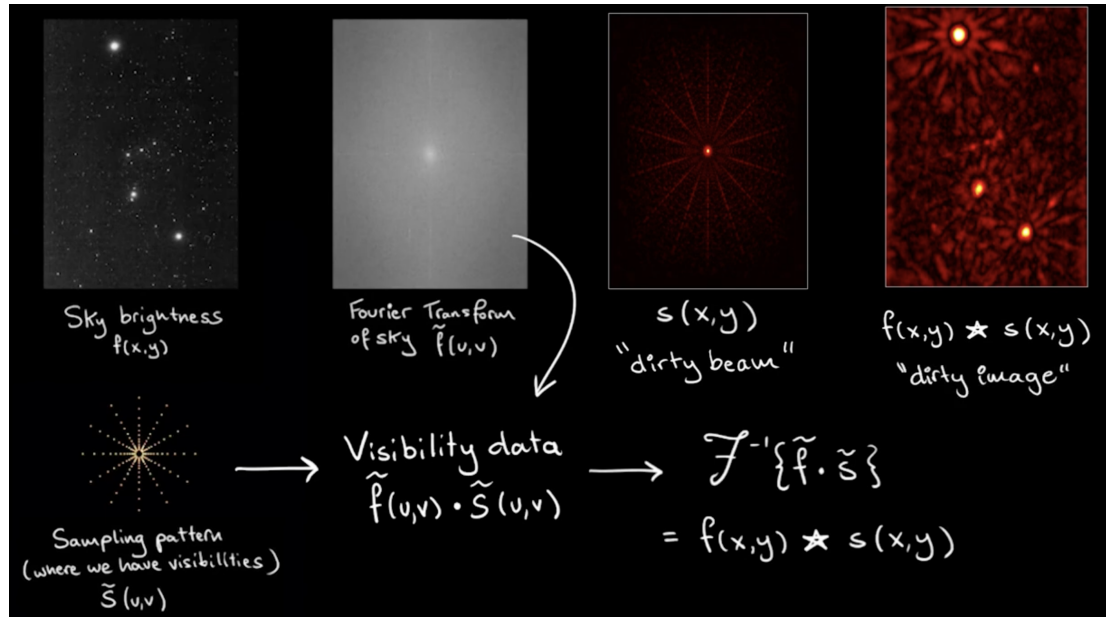


FIGURE 3.10: Example of an ALMA "dirty image". Taken from the "ALMA Primer Series: Introduction to CLEAN". Available at <https://www.youtube.com/watch?v=ff3KetPUyFE>.

3.5 ALMA image reconstruction using the CLEAN algorithm

As there is incomplete coverage of the u - v plane (e.g. Fig. 3.7) directly applying an inverse Fourier transform (iFT) on the visibility data cannot be used to obtain an image of the on-sky brightness distribution of the target source. Doing so, would reproduce a so called "dirty image" where the sky brightness distribution is convolved with the point spread function (PSF) or "dirty-beam" of the interferometer (see Fig. 3.10 for an example).

As such, a method of image reconstruction via aperture synthesis with interferometric visibility data was developed by (Högbom, 1974). This technique, called CLEANing, is outlined in Fig. 3.11 and is discussed in detail below in terms of its practical application to ALMA visibility data.

The CLEAN algorithm consists of two nested loops or cycles, an outer loop of "major cycles" and an inner loop of "minor cycles". The major cycle applies the FT and iFT to transform the data from the image domain into the Fourier domain (u - v plane) and

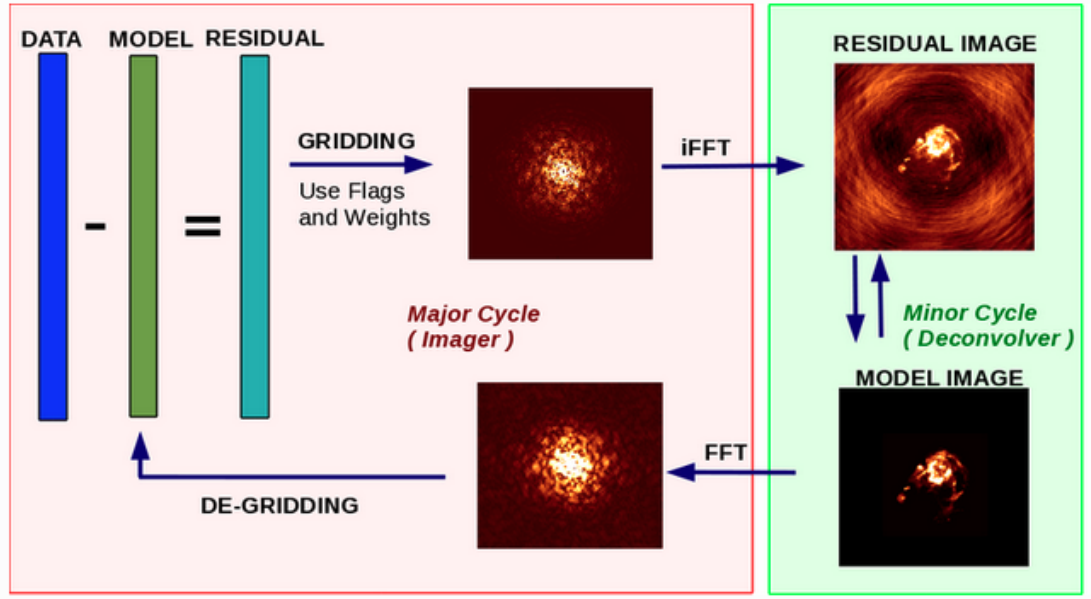


FIGURE 3.11: A schematic overview of the CLEAN algorithm. Taken from https://casadocs.readthedocs.io/en/stable/notebooks/synthesis_imaging.html. It is noted here that the large ring like structure in the residual image presented in this figure is due to the effect of incomplete sampling of the U-V plane.

vice versa, while, the minor cycle operates in the image domain only. The two cycles work together to perform an iterative and weighted χ^2 minimisation that solves Equation 3.19. In this equation $\vec{V}$ represents the complex visibilities, $\vec{I}$ is a list of parameters that model the sky brightness distribution (e.g. an image) and $[A]$ is some operator given by Equation 3.20 where $[F]$ is the 2D Fourier transform and $[S_{dd}]$ is the 2D spatial frequency sampling function. $[S_{dd}]$ is defined based on the observing setup.

$$\vec{V} = [A]\vec{I} \quad (3.19)$$

$$[A] = [F][S_{dd}] \quad (3.20)$$

Initially, the visibility data from an ALMA observation is transformed via an inverse fast Fourier transform (iFFT) into the image plane to begin the first "minor cycle". In the first minor cycle the algorithm identifies the brightest point in the image. This point is then added to a "model image" of the sky brightness distribution. The amplitude of this point is the brightness of the transformed image at this point times

a gain factor (less than one). This point on the model image is then called the first "clean component". This clean component is then convolved with the dirty beam and the model is transformed back into the u-v plane via an fast Fourier transform (FFT). The first major cycle then substitutes the modelled visibilities (from the transformed model image) from the visibilities observed by the ALMA array to produce "residual visibilities". These residual visibilities are transformed back into the image plane via an iFFT to begin the next minor cycle and the algorithm searches for the next clean component.

An issue with repeating the steps outlined above is that transforming from the image plane to the u-v plane via an FFT and back again via an iFFT is extremely expensive computationally. Instead, after the first clean component is added to the model, the model (convolved with the dirty beam) is subtracted from the initial residual image in the **image plane**. The algorithm repeats the minor cycle a few times adding new clean components, convolving with the dirty beam and subtracting from the residual image. After a few clean components are added an FFT is applied to go back into the u-v plane where the major cycle is repeated, new residual visibilities are calculated and then an iFFT is applied and a new residual image is generated. From here, the minor cycle can be repeated several times again and the whole process becomes cyclic until a stopping threshold is reached. This threshold is usually set by the user and can take a number of different forms such as a maximum number of major cycle iterations, a specified minimum value of flux density, or a specified number of standard deviations below the background root mean square noise (rms), below which, the CLEANing stops. In general, the ratio of major cycles to minor cycles gives an indication of the trade off between model accuracy and computational intensity (i.e the more major cycles the more accurate the residual image and model but at the cost of being more computationally expensive). The final cleaned image of the on sky brightness distribution is then produced by convolving the iterated model image of the on sky brightness distribution with a "clean beam" (Gaussian fit to the central part of the dirty beam) and adding on the final residual image.

3.6 Image weighting, image size and cell size.

As noted in the previous section and shown in Fig. 3.11, the CLEANing algorithm is a **weighted** χ^2 minimisation. As the coverage of points in the u-v plane is not an even distribution, the u-v plane must be divided into a grid of cells with a weight assigned to each cell so that areas that are well sampled in the u-v plane are not over-represented in the final reconstructed image. For example, in ALMA data, the u-v plane is usually better sampled towards the centre than further out which means structure at larger angular scales are better sampled and may be over-represented in the final image. As such, the weighting scheme chosen for the u-v plane effects the properties of the final extracted images. This section, provides a discussion of the different weighting schemes and the practical considerations when gridding the u-v plane.

3.6.1 Weighting schemes

There are three major weighting schemes, uniform, natural and robust/Briggs weighting. Figure 3.12 presents a schematic representation of uniform versus natural weighting scheme in a u-v plane. In uniform weighting, every cell in the u-v grid is assigned the same weight. Uniform weighting has the advantage of producing an image with high angular resolution, with the effect of beam side-lobes minimised. A disadvantage of this weighting method is that it produces a final image that is noisier (lower sensitivity) compared to the natural weighing scheme, hence making it more difficult to see extended structure in the final image.

In the natural weighting scheme, all cells are assigned a weight according the rms noise in each cell (making the grid weights look like a 2D Gaussian). The advantage of this weighting scheme is increased sensitivity and a higher signal-to-noise ratio. However, the disadvantage of this weighting scheme is that short baselines with better u-v coverage are more heavily emphasised resulting in an image with lower angular resolution (and a bigger synthesised beam) than compared to an image generated with a uniformly weighted grid. The last weighting scheme, Briggs (or robust) weighting

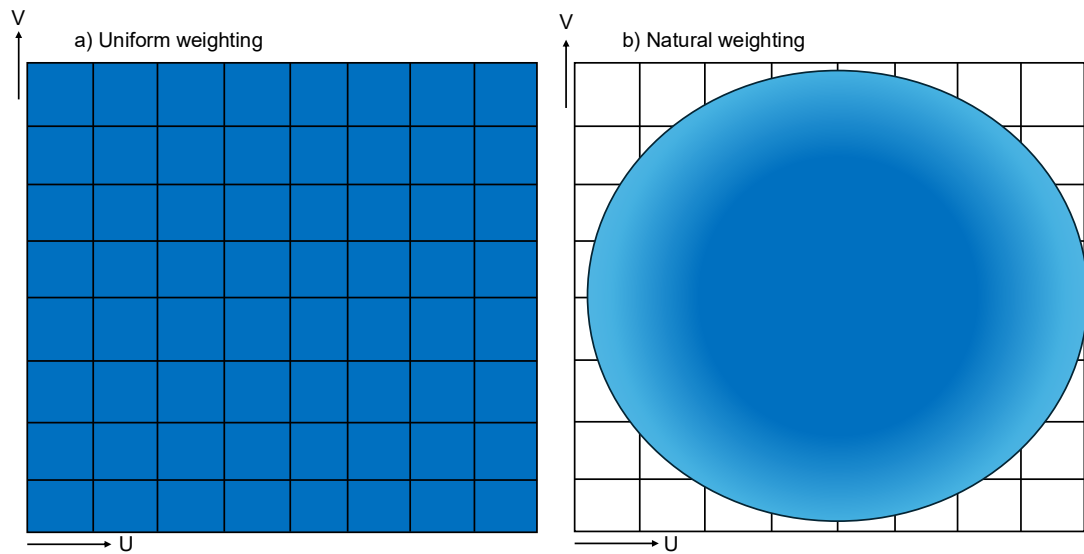


FIGURE 3.12: Schematic of uniform and natural grid weightings. Left: Uniform weighting scheme. Right: Natural weighting scheme.

is effectively a compromise between uniform and natural weighting (Briggs, 1995). Figure 3.13 summarises the idea behind Briggs weighting.

3.6.2 Image size and cell size

The size of the cells that divide the u-v plane (as shown in Fig. 3.12) can be chosen by the radio astronomer as part of the CLEANing process. The process of dividing the u-v plane into a grid of cells to apply weights to is called gridding. The size of the grid and the size of each cell in the grid is defined by two parameters: the image size and the cell size. This subsection provides some of the practical considerations to take into account when choosing suitable values for these parameters.

Due to the relationship between the image plane and the u-v plane, the FT scaling in one plane is related to the scaling in the other plane (see Fig. 3.14 and Equations 3.21 and 3.22). As such, both the image and u-v planes will always have the same number of cells.

$$\text{sidelength} = \text{cell size} \times \text{number of cells} \quad (3.21)$$

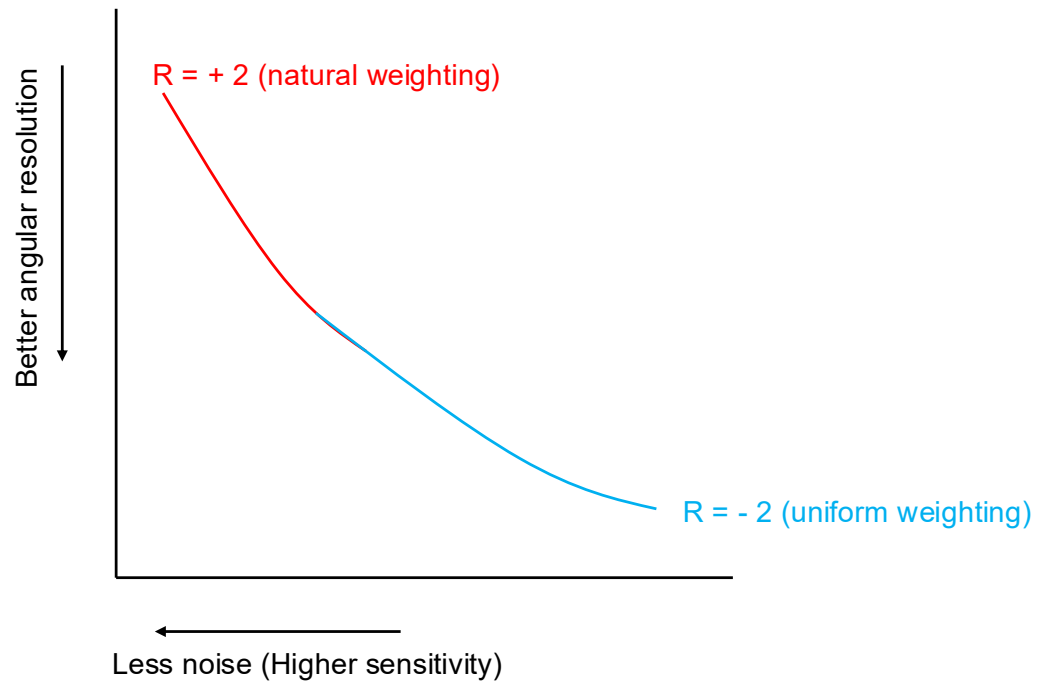


FIGURE 3.13: Briggs weighting.

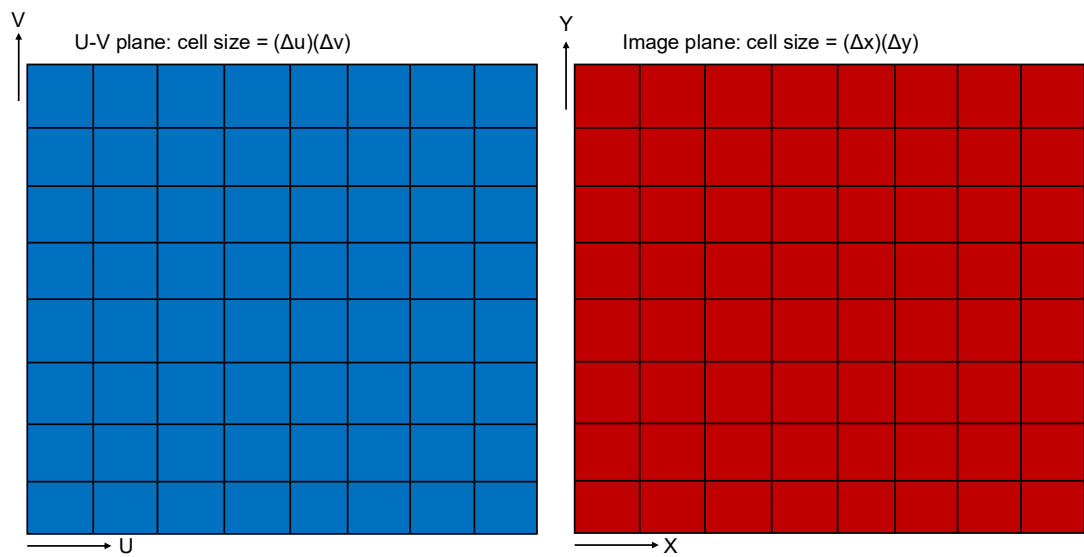


FIGURE 3.14: Schematic of u-v plane versus image plane gridding.

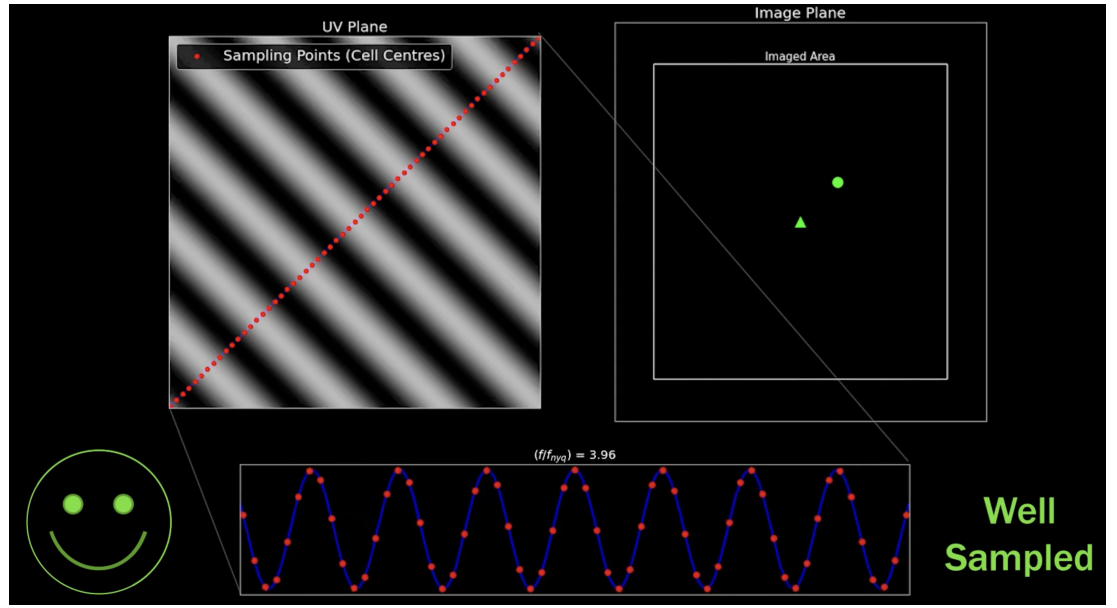


FIGURE 3.15: Example of good cell size sampling of the sinusoidal patterns in the $u-v$ plane. Taken from the "ALMA Primer Series: Image and Cell Size". Available at <https://www.youtube.com/watch?v=OC3IWpRRtEQ>.

$$\text{image cell size} \propto \frac{1}{u-v \text{ plane sidelength}}; \quad u-v \text{ plane cell size} \propto \frac{1}{\text{image plane sidelength}} \quad (3.22)$$

For example, for two point sources on the image plane (see Fig. 3.15, right). Intuitively, a cell size for the image plane that is small enough to resolve each point source could be selected. However, the $u-v$ plane must be considered (Fig. 3.15, left). Each pair of antennas in the interferometric observation produces a sinusoidal like pattern in the $u-v$ plane with unique amplitude and phase information. Given this, the choose a cell size must adequately sample this array of patterns in the $u-v$ plane. (i.e. the choosen cell size must at least Nyquist sample the highest frequency sinusoids in the $u-v$ plane.) In practice, this means that the cell size that is chosen should be small enough so that at least 6 to 8 cells cover the narrowest part of the synthesised beam (see Fig. 3.16 for an example).

The image size selected should be big enough to at least cover the field of view (FoV). For example, for a single antenna, the FoV is defined as the FWHM of the antenna

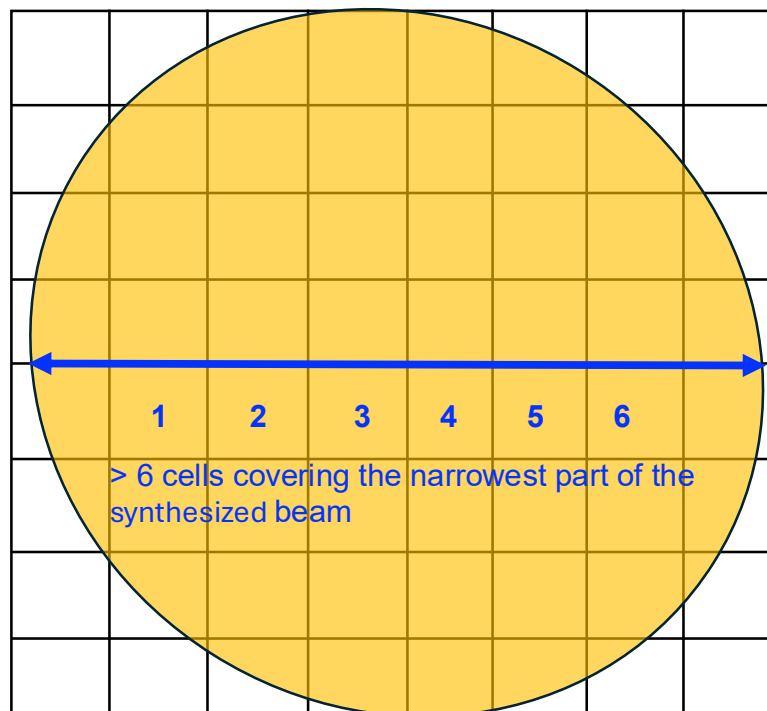


FIGURE 3.16: Schematic showing how the cell size well samples the synthesised beam,

response profile. Ignoring the minimal contributions of the beam side lobes to the antenna response pattern, the FoV is essentially the FWHM of the main beam (also known as the half power beam width (HPBW), see Fig. 3.2 and 3.17).

3.7 Working with ALMA data cubes

The end product interferometric observations with ALMA in dual polarization mode is a spectral data cube with flux density values at each pixel. The data cube has two spatial axes and a frequency (or equivalently, a wavelength axis or a velocity axis; provided a rest frequency is specified. These cubes are produced in CASA as the output of the CLEANing procedure in the `tclean` task). Figure 3.18 shows a schematic representation of the data cube structure.

From here, the data cubes are manipulated using the Cube Analysis and Rendering

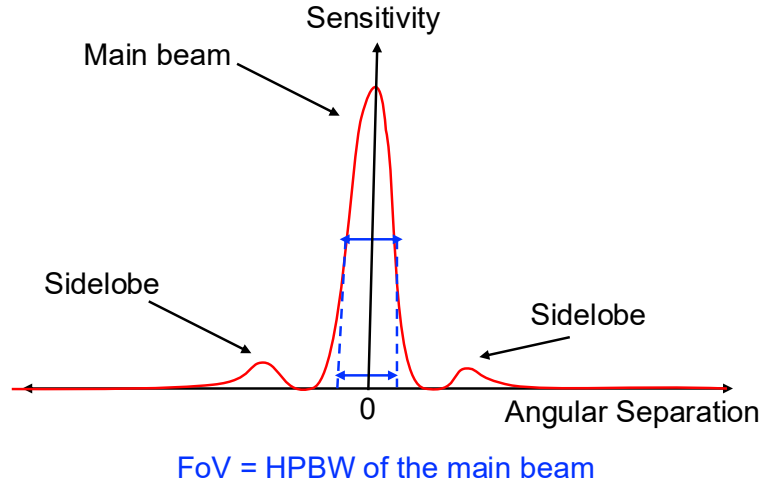


FIGURE 3.17: Schematic of 1D antenna response profile showing the FoV.

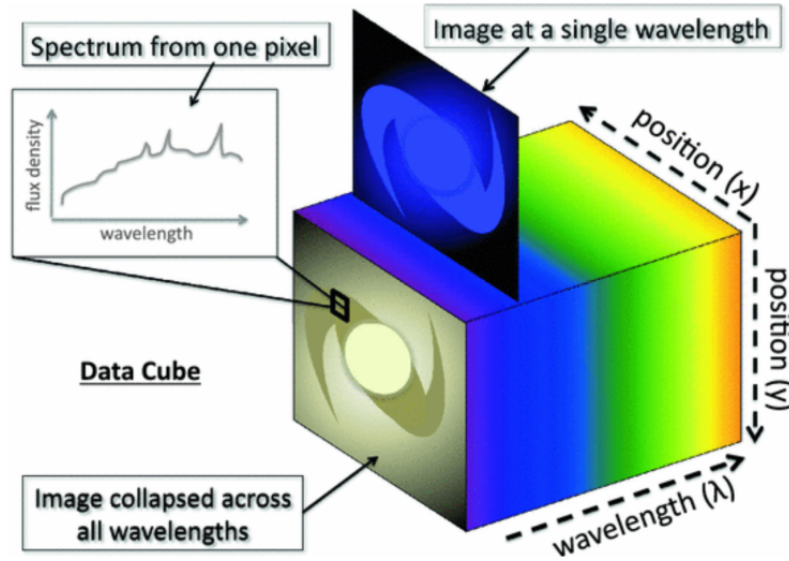


FIGURE 3.18: Schematic diagram of a spectral data cube. Taken from the online webpage of the PhD thesis entitled "Unlocking Galaxy Growth: A tale of two components".

Tool for Astronomy (CARTA) software. CARTA was used to generate the spectra, moment maps, and position-velocity (PV) diagrams presented in Chapter 6. From the files generated in CARTA, custom *Python* scripts were developed during the course of this thesis and used to apply data clipping, masking etc to tidy the output of CARTA

and produce the final figures presented in Chapter 6. The following subsections outline the methods used in CARTA and *Python* to produce the figures presented in Chapter 6.

3.7.1 Morphology: Channel Maps & Moment Maps

Channel maps are essentially an image at a single wavelength (or velocity) in the image cube. The channel maps were produced directly from the image cubes using the *Python* package *Go Fish* (Teague, 2019). This package reads in a cube of two spatial axes and a spectral axis (in velocity space). The package then uses file header information to convert the right ascension (RA) and declination (Dec) into an offset from the image centre in units of arcseconds. The user can specify a FoV and a velocity range to display. Figure 3.19 shows an example of a channel map extracted from an ALMA spectral image cube using *Go Fish*.

Moment maps are another tool that is useful for the visualisation and interpretation of ALMA spectral data cubes. The mathematical expressions that define moment maps are given in Equations 3.23 through 3.25

$$M_0 = \int I_\nu d\nu \quad (3.23)$$

$$M_1 = \frac{\int \nu I_\nu d\nu}{\int I_\nu d\nu} = \frac{\int \nu I_\nu d\nu}{M_0} \quad (3.24)$$

Higher order moments ($N \geq 2$) are defined as:

$$M_N = \frac{\int I_\nu (\nu - M_1)^N d\nu}{M_0} \quad (3.25)$$

In practice, the moment maps are generated using CARTA via the moments interface (Fig. 3.20). In this interface, the user selects the image cube to generate the moment of a spatial and spectral region of the cube to perform the calculation on, and whether

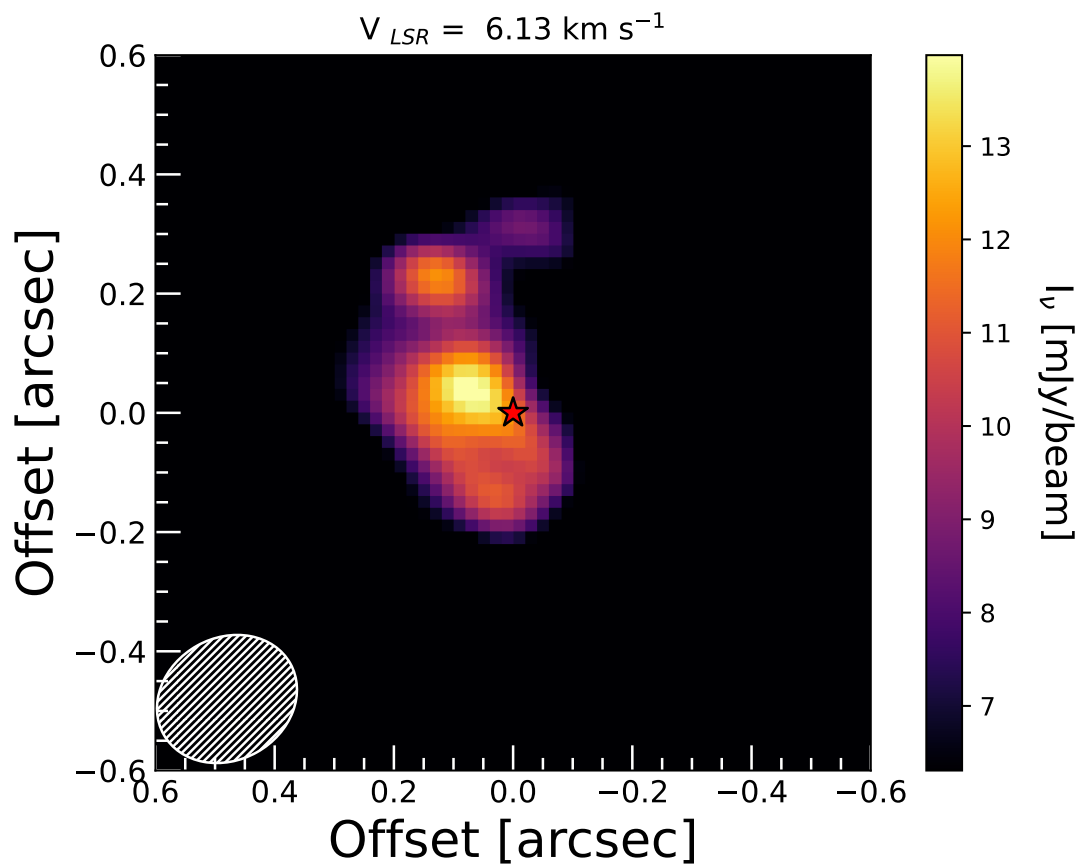


FIGURE 3.19: Channel map extracted from the SO 6_5-5_4 cube (see Chapter 6) using *Go Fish*. The source position is denoted by a red star, which is surrounded by gas emission traced by a molecular emission line.

or not to apply a mask to the pixels based on the value of flux density. CARTA allows for moment maps to be made from moment -1 to moment 11. The definition of each moment is given in Table 3.1. An example of a moment 0 (integrated intensity) map is presented in Fig 3.21. Following the production of the flexible image transport system (FITS) file containing the moment map in CARTA, a simple *Python* script was developed and used to apply data masking, clipping, and used to produce a tidy final version of figures presented in Chapter 6.

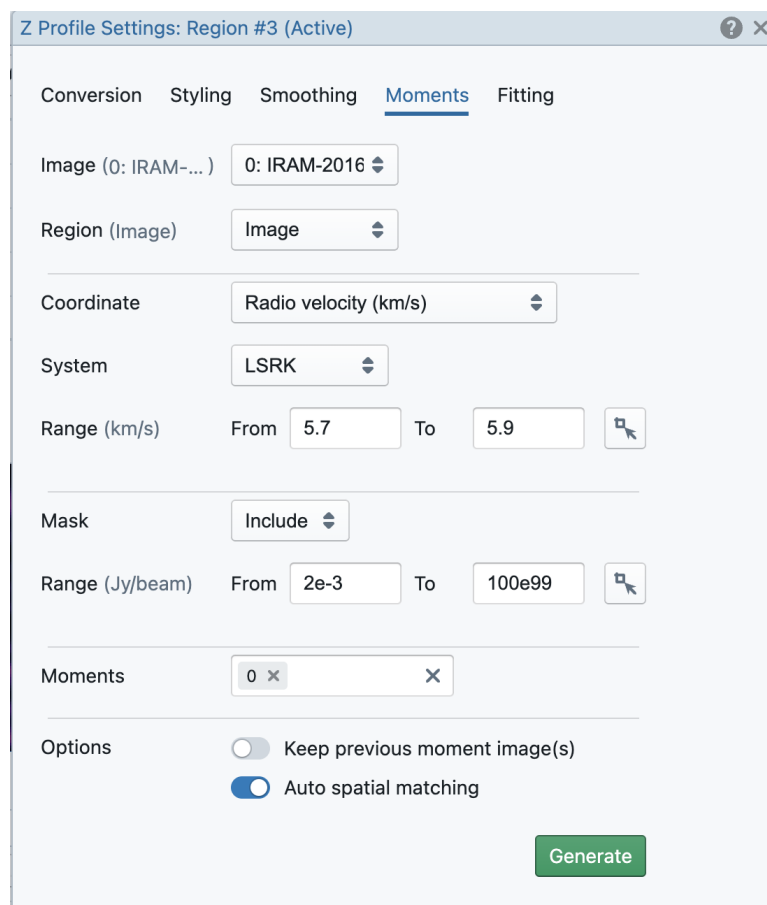


FIGURE 3.20: Window in CARTA for selecting parameters for generating moment maps.

3.7.2 Kinematics: Line Fitting & Position-Velocity diagrams

Line profiles and PV diagrams are useful tools to study the kinematics of the target object in an ALMA spectral cube. CARTA also has the functionality to extract these using

TABLE 3.1: Description of moments available in CARTA

Moment	Description
-1	Mean value
0	Integrated value
1	Intensity weighted coordinate/velocity field
2	Intensity weighted dispersion coordinate/velocity dispersion
3	Median value
4	Median coordinate
5	Standard deviation around the mean
6	Root mean square
7	Absolute mean deviation
8	Maximum value
9	Maximum coordinate
10	Minimum value
11	Minimum coordinate

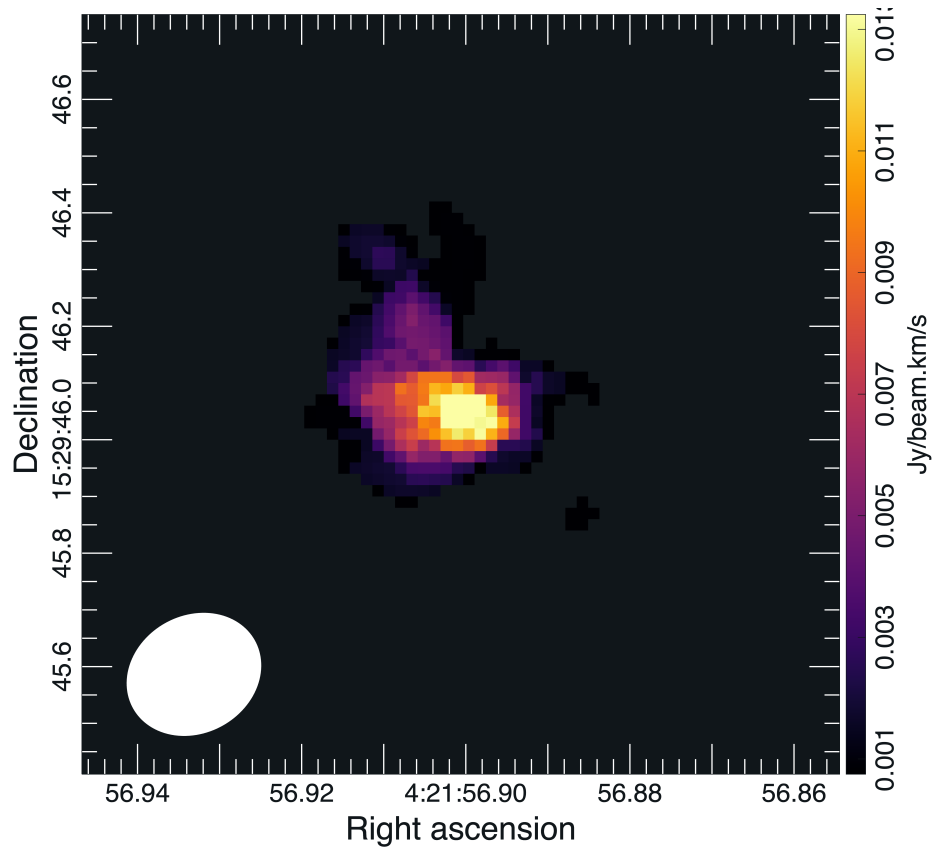


FIGURE 3.21: Example moment 0 map generated directly in CARTA. The image shown is the integrated intensity of the SO 65-54 molecular line emission integrated from 4.5 to 8.5 km s⁻¹. The synthesised beam is shown in the bottom left of the image.

user defined regions of the cube. In Chapter 6, the presented spectra are extracted initially via a circular aperture region centred at the source position. In Fig. 3.22, shows an example of a circular aperture placed on an image cube with the corresponding extracted flux spectrum. CARTA also has functionality to fit the extracted spectra using a non linear least squares (Levenberg-Marquardt) fitting method. The uncertainty on the fit is determined following the method outlined in Condon (1997). An example fit is shown in the left panel of this figure. The solid orange line represents the best fit. The orange dots represent the fit residuals.

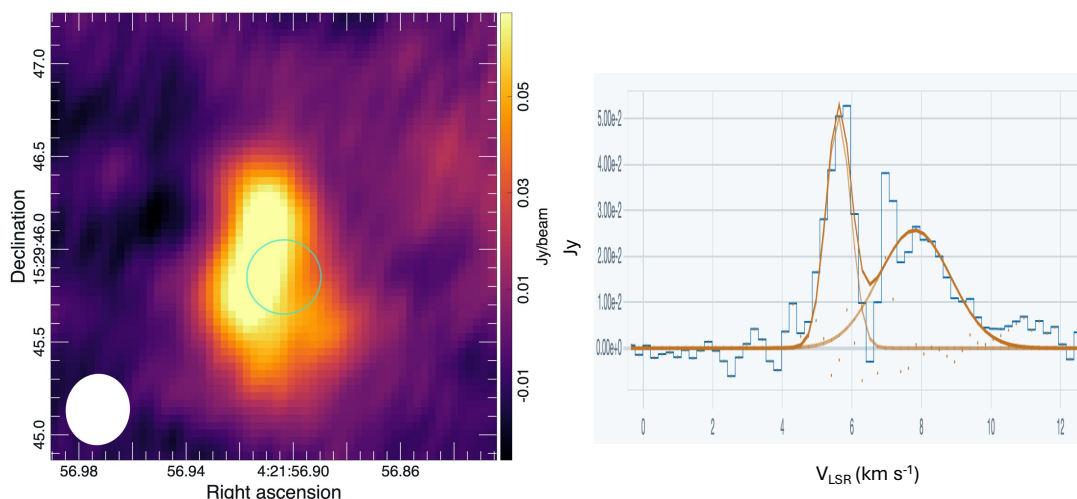
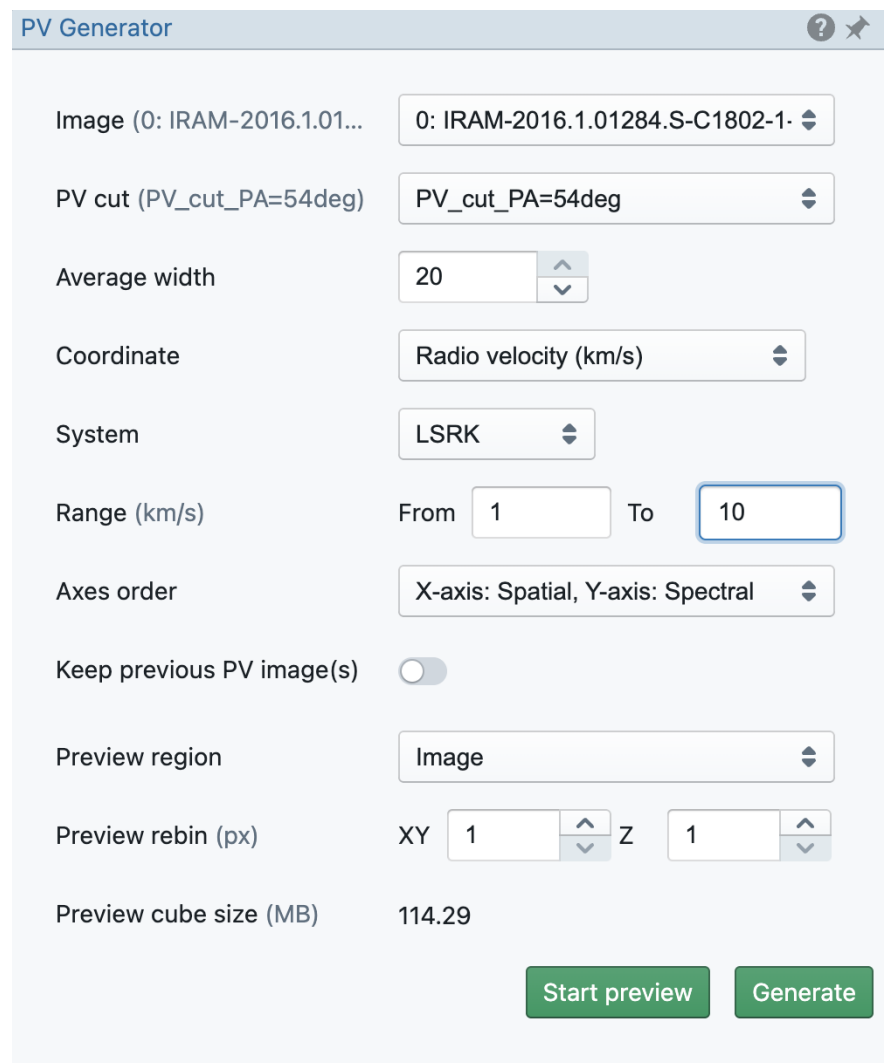


FIGURE 3.22: Example of a spectrum extraction in CARTA. On the left is a slice through the image cube with a cyan circle showing the spectrum extraction region. On the right is the corresponding spectrum extracted from the region marked with the cyan circle.

PV diagrams are extracted in a similar way in CARTA. In the version of CARTA used (4.1) in this thesis, there is no facility to extract a PV diagram via a 2D rectangular region. A line must be used rather than a two-dimensional pseudo-slit. The length and position angle of the line determines the length of the extraction slit and its position angle, whereas the width of the extraction slit is set by the user in terms of pixel units (i.e. how many pixels wide). An example of the PV extraction window in CARTA and an extracted PV diagram are shown in Figs. 3.23 and 3.24. Following the extraction of the PV diagrams a *Python* script is used to convert the colour map presented here into



The screenshot shows the 'PV Generator' window in the CARTA software. The window has a title bar with a question mark icon and a close button. The main area contains several configuration options:

- Image (0: IRAM-2016.1.01...)**: A dropdown menu showing '0: IRAM-2016.1.01284.S-C1802-1-...'.
- PV cut (PV_cut_PA=54deg)**: A dropdown menu showing 'PV_cut_PA=54deg'.
- Average width**: A numeric input field with '20' and up/down arrow buttons.
- Coordinate**: A dropdown menu showing 'Radio velocity (km/s)'.
- System**: A dropdown menu showing 'LSRK'.
- Range (km/s)**: Two numeric input fields labeled 'From' and 'To', with values '1' and '10' respectively.
- Axes order**: A dropdown menu showing 'X-axis: Spatial, Y-axis: Spectral'.
- Keep previous PV image(s)**: A toggle switch that is currently turned off.
- Preview region**: A dropdown menu showing 'Image'.
- Preview rebin (px)**: Two numeric input fields labeled 'XY' and 'Z', both with the value '1' and up/down arrow buttons.
- Preview cube size (MB)**: A text label followed by the value '114.29'.

At the bottom right of the window, there are two green buttons: 'Start preview' and 'Generate'.

FIGURE 3.23: PV diagram generator window in CARTA.

the contour maps like the ones presented in Chapter 6.

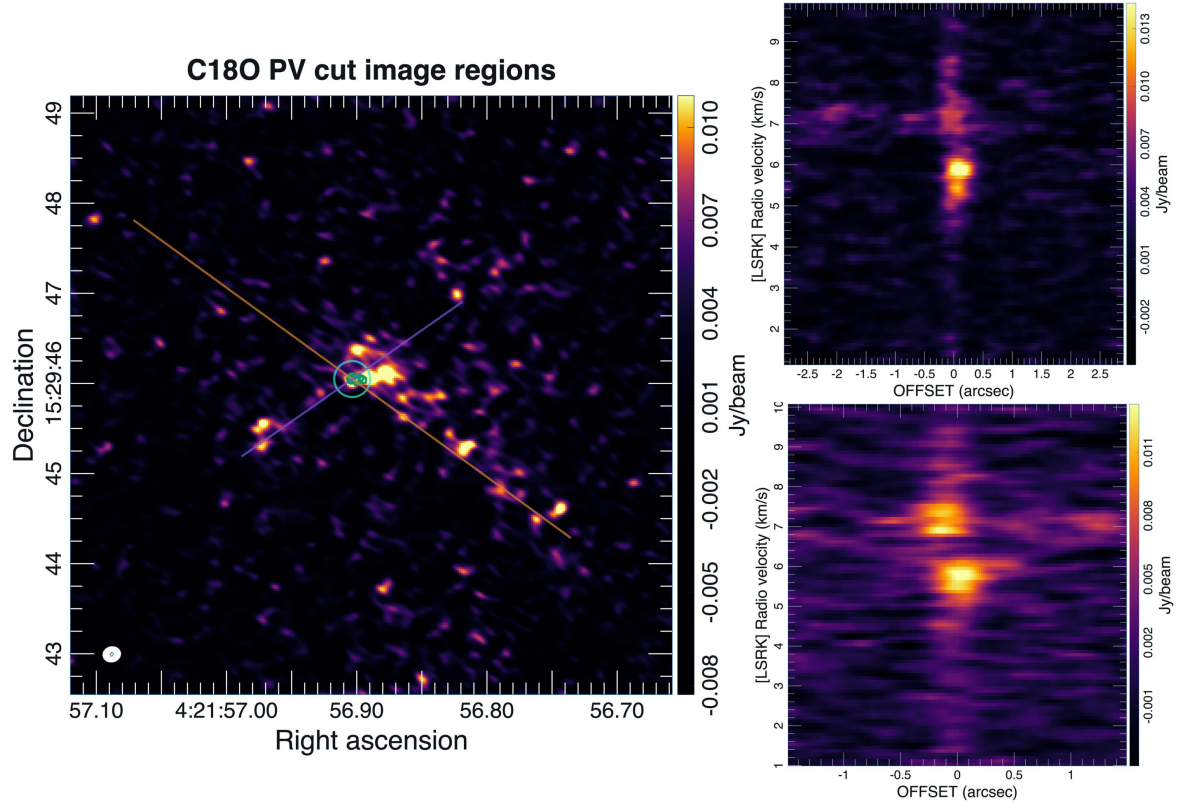


FIGURE 3.24: Example of PV diagrams directly exported from CARTA. Left: A cut through the C^{18}O (2–1) image cube with the PV extraction regions marked (the orange and purple lines). The cyan circle represents the region where a spectrum may be extracted. The green contours show the continuum emission (dust emission) in Band 6 (1.1 to 1.4 millimeters). The synthesised beams for the continuum and C^{18}O (2–1) are shown in the bottom left corner of this panel as the green and white ellipses respectively. Top Right: PV diagram extracted through the orange extraction region. Bottom Right: PV diagram extracted through the purple extraction region.

Chapter 4

Constraining the extent of the [O I] $\lambda 6300$ low velocity component emission line region with spectro-astrometry

4.1 Chapter Introduction

The goal of this chapter is to expand on the pilot study of Whelan et al. (2021) in an attempt to ascertain the origin of the FEL LVC and to test the assumption made by disc wind driven accretion models that the mass loss rate for the low velocity wind should be comparable to the stellar mass accretion rate. Prior to this pilot study, several previous works (e.g. Banzatti et al. (2019), Fang et al. (2018), and Simon et al. (2016)) attempted to establish the origin of the optical FEL LVC through studying the kinematics of (a combined total of) hundreds of optical FEL profiles from a vast sample of CTTs from various star forming regions (see Section 1.2.5). From these studies, a consensus was reached regarding the broad part of the LVC, the LVC-BC. This sub-component likely traces the base of an MHD disc wind (see Pascucci et al. (2020) and Chapter 1). However, the spatially compact nature of the LVC, presents a challenge to these previous kinematic studies. No spatial information regarding the LVC was extracted in these previous works and as such, the origin of the LVC-NC as either an MHD or photoevaporative wind remains elusive.

The work of Whelan et al. (2021) presented SA as a novel approach to understanding the optical FEL LVC as this technique can extract sub seeing spatial information about

the FEL emission from seeing limited observations. This pilot study presents spectro-astrometric analysis of the CTTS RU Lupi in both the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs, providing the first evidence that the LVC-NC may be tracing an MHD disc wind. In the case of RU Lupi, the LVC-NC was traced to within ~ 1 AU from the position of the star, meaning that the flow is launched well within the gravitational potential well of RU Lupi. Alongside this, a clear decrease in the velocity of the flow with increasing height above the disc mid-plane is observed for the LVC-NC of RU Lupi (see their Fig. 2). This negative velocity gradient is the opposite of what is expected for high velocity jets (Ferreira, 1997; Ferreira, Dougados, and Cabrit, 2006) and is indicative of a disc wind (Pontoppidan, Blake, and Smette, 2011). The combination of both a negative velocity gradient and a launch region within the potential well of RU Lupi clearly suggests an MHD wind origin for the LVC-NC.

However, there remains an open question about whether the optical FEL LVC regularly traces an MHD disc wind or if RU Lupi is a unique case. In using SA to extract spatial information about the LVC, the aim of this chapter is to test this and expand the work of Whelan et al. (2021) to a larger sample of CTTSs with varying physical properties. Additionally, this chapter aims to examine the assumption made by wind driven accretion models; namely, that the mass outflow rate of the low velocity wind should be comparable to the stellar mass accretion rate.

4.2 Sample, observations & methods

This section describes the sample of CTTSs selected and the methods used to analyse the data. The reduction and post processing steps applied to the UVES data has been described previously in Chapter 2 alongside a description of the spectro-astrometric technique and its application to UVES data.

4.2.1 Sample of targets

The sample of selected targets for this study consists of ten well known CTTSs with a range of physical properties as shown in Table 4.1. The sample covers a range of

disc inclination angles from close to face on (e.g. TW Hya) to almost edge on (e.g. GK Tau). The sample also covers a range of stellar masses (from 0.5 to 2.2 $M_{\odot}$) and contains a combination of high and low accretors. The sample also contains a known outbursting star (EX Lup) and three transition disc objects. These targets were selected on the basis of this diversity of properties. Variation in the mass of the selected targets is important as it is known that the stellar mass is correlated to the stellar mass accretion rate (e.g. Manara et al., 2023; Testi et al., 2022), which itself is thought to be linked to ejection processes (Cabrit, 2007; Cabrit et al., 2010). Therefore, testing the spatial and spectral properties of the LVC over this range of masses is crucial to potentially constrain its origin. Additionally, targets of varying disc inclination are selected to account for the variation in spatial properties of the LVC with inclination (i.e. to account for the effect of disc inclination on the recovered spatial properties of the LVC). By performing both kinematic fitting and spectro-astrometric analysis on this sample, it is therefore expected that this sample provides a robust testing ground to examine the LVC. The [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs were chosen for this kinematic and spectro-astrometric analysis as they are the brightest FELs with known detections of the LVC (e.g. McGinnis et al., 2018; Nisini et al., 2024; Whelan et al., 2021, etc.). The [O I] $\lambda 5577$ FEL was not included as for most of these targets (excluding DG Tau to be discussed in Chapter 5) this FEL is faint and only presents a HVC.

TABLE 4.1: Properties of each target star relevant to this study.

Source	$M_{\star}$ [$M_{\odot}$]	$R_{\star}$ [$R_{\odot}$]	Dist. ^a [pc]	SpTy	Disc Type	Disc PA ^b [$^{\circ}$]	Disc Incl. [$^{\circ}$]	A_v [mag]	Drives Jet?	Refs.
TW Hya	0.61	1.20	60.14 ± 0.06	K7	transition	332	7	0.0	No	C11,H08,P08,P&S09,F18,F06
T Cha	1.50	1.80	102.72 ± 0.35	G8	transition	113	69	1.7	No	Ca19,Hu18,P&S09,S14,Sc09
MP Mus	1.20	1.40	97.89 ± 0.12	K1IVe	full	167	32	0.7	No	C11,Cr09,S14,P20,
LkH α 330	2.20	1.50	317.94 ± 3.45	G3	transition	165	12/42 ^c	1.8	No	Wi16,I13
GQ Lup	0.89	1.80	154.10 ± 0.68	K7	full	346	60	0.5	No	JK13,D12,Mc12,F18,B19
DG Tau	0.80	1.85	125.26 ± 1.83	K6	full	136	32	1.6	Yes	B19,G16,M14,Fu06,Hi97
CW Tau	1.4	1.94	131.55 ± 0.64	K3	full	54	65	2.8	Yes	B19,Co12,Po11,Fu06,HI97
GK Tau	0.69	1.80	129.14 ± 0.30	K7	full	300	71	0.9	Yes	B19,F18,Si16,GH98
DK Tau	0.68	1.84	132.02 ± 0.88	K7	full	4	26	1.3	Yes	B19,F18,GH98,Fu06
EX Lup	0.50	1.60	154.72 ± 0.36	M0.5	full	63	38	1.1	Yes ^d	B19,F18,A17

A17 = Alcalá et al. (2017); B19 = Banzatti et al. (2019); Ca19 = Cahill et al. (2019); C11 = Curran et al. (2011); Co12 = Coffey et al. (2012); Cr09 = Cortes et al. (2009); D12 = Donati et al. (2012); F18 = Fang et al. (2018); Fu06 = Furlan et al. (2006); G16 = Grankin (2016); GH98 = Gullbring et al. (1998); H08 = Herczeg and Hillenbrand (2008); Hu18 = “The Disk Substructures at High Angular Resolution Project (DSHARP). II. Characteristics of Annular Substructures” (2018); I13 = Isella et al. (2013); JK13 = Johns-Krull et al. (2013); K18 = Kurtovic et al. (2018); M14 = Maurri et al. (2014); Mc12 = McClure et al. (2012); P08 = Pascucci et al. (2008); P&S09 = Pascucci and Sterzik (2009); P20 = Pascucci et al. (2020); Po11 = Pontoppidan, Blake, and Smette (2011); S14 = Sacco et al. (2014); Sa13 = Salyk et al. (2013); Sc09 = Schisano et al. (2009); Si16 = Simon et al. (2016); Wi16 = Willson et al. (2016);, Hi97 = Hirth, Mundt, and Solf (1997).

^a = The distances to each target and their corresponding uncertainties were calculated using the parallax values and uncertainties from Gaia data release 3 (Prusti et al., 2016; Vallenari et al., 2023). The distance to each target (in parsecs) is taken to be the inverse of the measured Gaia parallax (in milliarcseconds)

^b = PA of the disc major axis.

^c = This source has split disc with an inner and outer part two different inclination angles. I use the outer inclination angle of 42° for the length and velocity deprojection described later in the chapter.

^d = The jet has been detected through spectro-astrometric analysis and will be presented in a future paper by Unni et al. (in prep).

4.2.2 Observations & data reduction

The spectra for this sample of targets were obtained using the UVES instrument (see Chapter 2 and Dekker et al. (2000) for more information on the instrument.) on the VLT. Spectra were taken from May to June 2012 via programme 089.C-0299(B) (P.I. I. Pascucci) and from October 2021 to April 2022 via programme 108.22M8.001 (P.I. E. Whelan). The wavelength range achieved was 5000 - 7000 Å with a spectral resolution of $R \sim 50,000$ ($\sim 6 \text{ km s}^{-1}$). The slit width was set to 1" for all sources except for MP Mus (observed on May 2 2012), where a 0".3 slit was used. The slit width was changed in this case as the seeing conditions had improved significantly. A narrower slit is therefore required in order for the slit to maintain even illumination during observation to avoid spatial artefacts in the spectro-astrometric analysis (Brannigan et al., 2006). The observational log for these two programmes is presented in Table A.1 of Appendix A.

These data were initially reduced using the UVES data reduction pipeline (version 6.1.6; Freudling et al., 2013b) within the *EsoReflex* environment to produce both 1D and 2D spectra. As discussed in Chapter 2, after the initial pipeline reduction, the 2D spectra were cleaned of skylines, corrected for cosmic ray effected pixels and were continuum subtracted using a combination of IRAF tasks and a custom *Python* script. The 1D spectra of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ emission lines were obtained from the UVES pipeline by collapsing the 2D spectra along the full width of the slit. These line profiles corrected for contaminating telluric and photospheric features using ESOs' MOLECFIT pipeline (Kausch et al., 2015; Smette et al., 2015) and the PRT available in the STAR MELT software (Campbell-White et al., 2021) which essentially follows the method presented by Hartigan et al., 1989. Table 4.2 shows the list of the WTTS templates used to remove the contamination photospheric absorption features present in the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ lines of the target CTTSs in the sample. Best practice dictates that the spectral type of the target CTTS and the template WTTS be as close as possible in order to attain a good fit to the present photospheric features during the removal process (Stock et al., 2020, 2022). In most cases the spectral type of the WTTS

TABLE 4.2: WTTS templates used for photospheric correction of the target stars in the sample.

Target	SpTy	WTTS template	SpTy
TW Hya	K7	HIP17157	K7
T Cha	G8	HD69830	G8
MP Mus	K1IVe	HD114783	K1
LkH α 330	G3	V1077 Tau	G8
GQ Lup	K7	HIP17157	K7
DG Tau	K6	2MASS J15430624-3920194	K6
CW Tau	K3	HD99492	K3
GK Tau	K7	MW Lup	K7
DK Tau	K7	MW Lup	K7
EX Lup	M0.5	Gl846	M0.5

template matched the spectral type of the target CTTS. In some cases (e.g. MP Mus) a WTTS template of the targets exact spectral type was unavailable in STAR MELT so, a WTTS template of as close as possible spectral type was selected as a substitute. Regardless, a suitable fit was obtained with minimal regions of over/under subtraction of the contaminant features.

4.2.3 Analysis methods: SA and kinematic fitting

As discussed in Section 2.2.2 of Chapter 2, the [O I] λ 6300 and [S II] λ 6731 line profiles are shifted into the stellocentric frame from the topocentric frame after a radial velocity and heliocentric velocity correction had been applied to each target CTTS spectrum. The radial velocity was calculated with both the Li λ 6707 and Ca λ 6439 absorption lines. This was done as it has been found that solely relying on the Li λ 6707 absorption line for finding a radial velocity may introduce unwanted uncertainties given that the central wavelength of this absorption line appears to be correlated to the effective temperature of the CTTS (Campbell-White et al., 2023). After each of the line profiles are shifted into the stellocentric frame, kinematic fitting was performed to decompose each line into its constituent sub-components. The method outlined in Chou et al. (2025) was used to decompose each line profile. This is a χ^2 minimisation method which involves fitting an increasing number of Gaussian components to the target line profile until the χ^2 statistic is **not** reduced by at least 20% compared to the

χ^2 value calculated from a previous model, which contained one fewer Gaussian sub-components. Figure 4.1 shows an example of this kinematic fitting procedure used to reproduce the [O I] $\lambda 6300$ line of CW Tau. Each Gaussian sub-component in this figure is a solid coloured line, with the overall fit to the observed line profile marked as a black dashed line. As is observed from Fig. 4.1, as the number of individual Gaussian components increases, there is a significant reduction in the value of the χ^2 statistic, indicating a better fit of the model fit to the observed data. The uncertainty on the peak velocity and FWHM of each fitted sub-component is determined via the method of Nisini et al. (2024) (as summarised by Equation 2.8, in this context, S/N represents the signal-to-noise in each sub-component). This method was chosen so that the kinematic fitting presented in this chapter could be compared to previous kinematical studies that have used similar χ^2 minimisation methods for line profile decomposition (e.g. Banzatti et al., 2019; Chou et al., 2025; Fang et al., 2018; Nisini et al., 2024).

However, for all CTTSs in the sample the uncertainty on the stellar radial velocity calculated for each target CTTS is larger than the uncertainty on the fitted peak velocity and FWHM of each sub-component using the Nisini et al. (2024) method. As such, the uncertainty on the peak velocities and FWHM vales of each sub-component was taken as the minimum uncertainty on the radial velocity of each star (see Chapter 2, Section 2.2.2 for details regarding the radial velocity calculations and uncertainties).

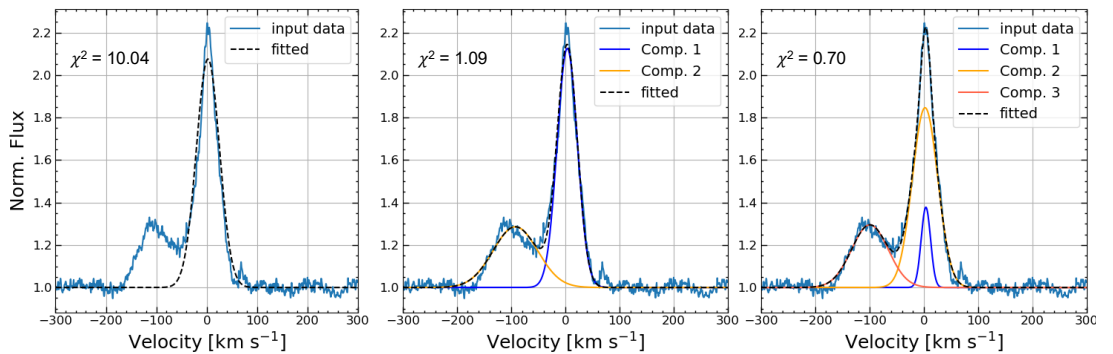


FIGURE 4.1: Example of the kinematic fitting procedure outlined by Chou et al. (2025). Shown is the continuum normalised [O I] $\lambda 6300$ line profile of CW Tau with one (left panel), two (middle panel) and three Gaussian sub-components (right panel) with the corresponding χ^2 values calculated for each fit.

The spectro-astrometric analysis was performed as described in Section 2.1 of Chapter 2. Each of the spatial profiles extracted along each emission line region was fit with a Gaussian function, from which, the centroid and FWHM of each spatial profile were measured. Position spectra and Δ FWHM spectra for each target/line were then constructed as described in Chapter 2.

In order to recover the actual extent and size of the emission as traced by SA, each 2D spectrum was continuum subtracted and each spatial profile covering the emission line region was refit as described in Section 2.1 of Chapter 2. The uncertainty on each measured centroid was calculated using Equation 2.4. The value quoted for each position spectrum presented in this chapter before/after continuum subtraction is the average centroid uncertainty calculated over the region containing the FEL before/after continuum subtraction.

To determine if the centroid offset in the position spectrum is real; repeated fitting of the spatial profiles from anti-parallel slit PAs are used to determine if the offset measured "flips" as the slit is rotated through 180° (see Chapter 2 and Whelan and Garcia (2008)). Figure 4.2 and 4.3 shows an example of the spectro-astrometric analysis of the [O I] $\lambda 6300$ line of TW Hya in all four slit PAs, two along the outflow axis and two along the disc axis. The top panels of each figure is the continuum normalised [O I] $\lambda 6300$ line profile at each PA. The centre panels show the position spectra of the [O I] $\lambda 6300$ at each PA with the continuum subtracted centroids over-plotted. Lastly, the bottom panels are the Δ FWHM spectra of the [O I] $\lambda 6300$ at each PA with the continuum subtracted centroids over-plotted. A true spectro-astrometric detection is defined as a signal in the position spectrum measured above 3σ (σ determined via Equation 2.4) that "flips" when the slit is rotated through 180° (Whelan and Garcia, 2008; Whelan et al., 2021; Whelan et al., 2024). From Fig. 4.2, it is clear that no such "flip" in the position spectrum is observed when the slit is rotated through 180° for the jet axis PAs. Similarly, no "flip" is observed in the signal is observed in the disc PAs. What

is concluded from this is that the offset measured from TW Hya in this line isn't a true spectro-astrometric offset and as such, the average centroid uncertainty is taken as lower limit on the extent of the emission traced by this FEL above the disc mid-plane (i.e., X_c). The wind height above the disc midplane, l_{wind} , is then taken as $3 \times X_c$ deprojected to account for the disc inclination. These non detections in the position spectra were the case for the majority of the CTTs in the sample as discussed later in Section 4.4.

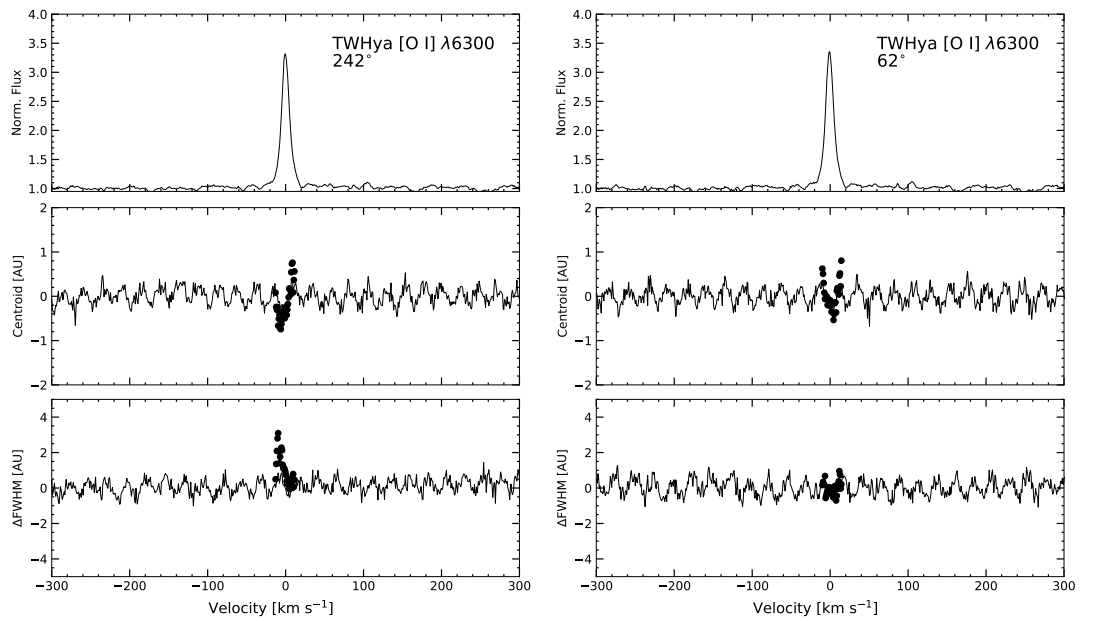


FIGURE 4.2: Example showing the spectro-astrometric analysis of TW Hya for the [O I] $\lambda 6300$ line along the jet PAs. No "flip" is observed for the measured signal in the position spectra as the slit is rotated through 180° . This is expected if the signal is real (Whelan and Garcia, 2008).

It is noted here that in both Figs. 4.2 and 4.3, there appears to be a repeating pattern in the non continuum subtracted position spectrum (solid black line in the middle panels of these figures). This is a spatial aliasing effect, that is caused by the resampling of the wavelength/velocity axis by the UVES pipeline during the initial data reduction. This is shown by Fig. 4.4. This figure shows the same spectro-astrometric analysis performed on the 152° PA of the [O I] $\lambda 6300$ line of TW Hya before wavelength calibration had been performed. Clearly, in this figure the repeating pattern is no longer present in the position of $\Delta FWHM$ spectra. Additionally, this aliasing effect has been

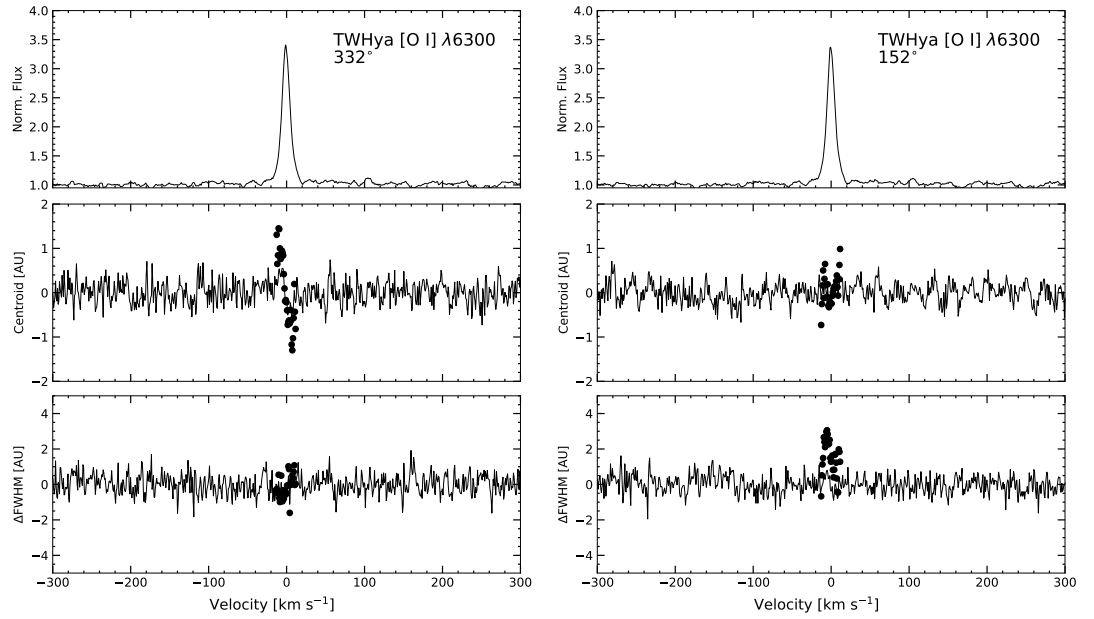


FIGURE 4.3: Example showing the spectro-astrometric analysis of TW Hya for the [O I] $\lambda 6300$ line along the disc PAs. No "flip" is observed for the measured signal in the position spectra as the slit is rotated through 180° . This is expected if the signal is real (Whelan and Garcia, 2008).

previously noted in the case of the X-Shooter instrument by Whelan et al. (2015). This aliasing effect only seems to appear for TW Hya and GQ Lup. Additionally, there is a large scatter in the continuum subtracted points (black points) in the position and ΔFWHM spectra at the wings of the line due to the decrease in the S/N ratio at the line wings. This limits the accuracy of the continuum subtraction at the line wings and increases the uncertainty on the SA (see Equation 2.4).

Lastly, the uncertainty on the ΔFWHM was measured as follows: As the *Python* package `scipy.optimize.curve_fit` was used to fit a Gaussian to each spatial profile, a covariance matrix is computed to estimate the "goodness" of the fit parameters. The square root of the diagonal entries in this covariance matrix correspond to the standard deviation on each fit parameter for the Gaussian used to fit each spatial profile (e.g. centre, amplitude etc). The uncertainty of the measured FWHM is taken from its corresponding entry in this covariance matrix and the values quoted in the ΔFWHM spectra presented in this chapter is the average centroid FWHM calculated over the region containing the FEL before/after continuum subtraction. These are converted

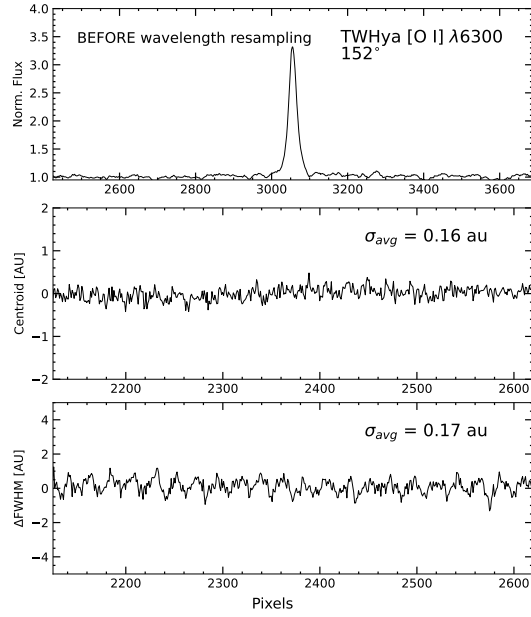


FIGURE 4.4: The spectro-astrometric analysis of TW Hya for the [O I] $\lambda 6300$ line at PA = 152° **before** wavelength axis resampling (i.e. the raw data with the dispersion axis in pixels). Note the original 1D and 2D arrays are of different lengths in pixels but they have been aligned to cover the equivalent of -300 to 300 km s^{-1} . It was observed that the repeating pattern previously seen in the centre right of Fig. 4.3 (i.e. the solid black line; before continuum subtraction) is no longer present (see centre panel of this figure). This indicates that the repeating patterns seen in Figs. 4.2 and 4.3 is an aliasing effect due to the wavelength axis resampling during the pipeline data reduction process.

from arcseconds to AU given the assumed distance of each target star (see Table 4.1).

4.3 Results

This section presents the results of the kinematic and spectro-astrometric analysis of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ forbidden emission lines present in the sample of ten CTTs observed during the two UVES programmes during periods 89 and 108 respectively.

4.3.1 Kinematic Fitting

Figs. 4.5 and 4.6 show the kinematic fitting of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ lines present in the spectra of the CTTs in the sample respectively. The spectra are taken from the PAs for observations along the known jet PA for each star. The flux of the

line profiles presented in these figures have been normalised with respect to the local continuum level around each line profile. Each line profile has been shifted into the stellocentric rest frame by applying the radial and heliocentric velocity corrections as outlined in Chapter 2. Each sub-component of the line profiles has been shaded in different colours and follows the same convention that has been used in previous studies, green for the HVC, red for the LVC-BC and blue for the LVC-NC (Fang et al., 2018, 2023a; Pascucci et al., 2020). The corresponding results of the kinematic fitting for both sets of line profiles are shown in Tables 4.3 and 4.4 for the [O I] $\lambda 6300$ and the [S II] $\lambda 6731$ FELs respectively. The [O I] $\lambda 6300$ and [S II] $\lambda 6731$ lines were selected for this kinematic and spectro-astrometric analysis as the critical density for each line ($\sim 2 \times 10^6$ for [O I] $\lambda 6300$ and $\sim 2 \times 10^4 \text{ cm}^{-3}$ for [S II] $\lambda 6731$; Whelan et al., 2021) implies that these lines trace different layers of the outflow structures seen in the CTTs in this sample (Chou et al., 2025; Giannini et al., 2019). Also, these lines have been fit with similar kinematic fitting analysis for many other CTTs, in many previous studies (e.g. Banzatti et al., 2019; Fang et al., 2018; Nisini et al., 2024). As such, comparison to these previous studies will aid in the interpretation of the analysis presented in this chapter.

The kinematic analysis of the [O I] $\lambda 6300$ show a clear separation between the stars in this sample. It is clear that the six of the ten [O I] $\lambda 6300$ line profiles are reproduced by fitting multiple Gaussian sub-components. These stars will be referred to as multi component (MuC) sources in this chapter for consistency with the existing literature. Conversely, four of the sources in this sample (TW Hya, T Cha, MP Mus and LkH α 330) have their [O I] $\lambda 6300$ line profiles reproduced with a single Gaussian component. Again, for consistency with the literature, these stars will be referred to as single component (SC) sources for this chapter.

Interestingly, of the [S II] $\lambda 6731$ line profiles presented in Fig. 4.6, none of the SC sources in the sample show a detectable [S II] $\lambda 6731$ line in their jet PA spectra. This may hint that the micro-jet emission these sources may be very compact as the [S II] $\lambda 6731$ should trace more extended outflow emission given its lower critical density compared to the [O I] $\lambda 6300$ FEL (Giannini et al., 2019; Whelan et al., 2021). Alongside this, no

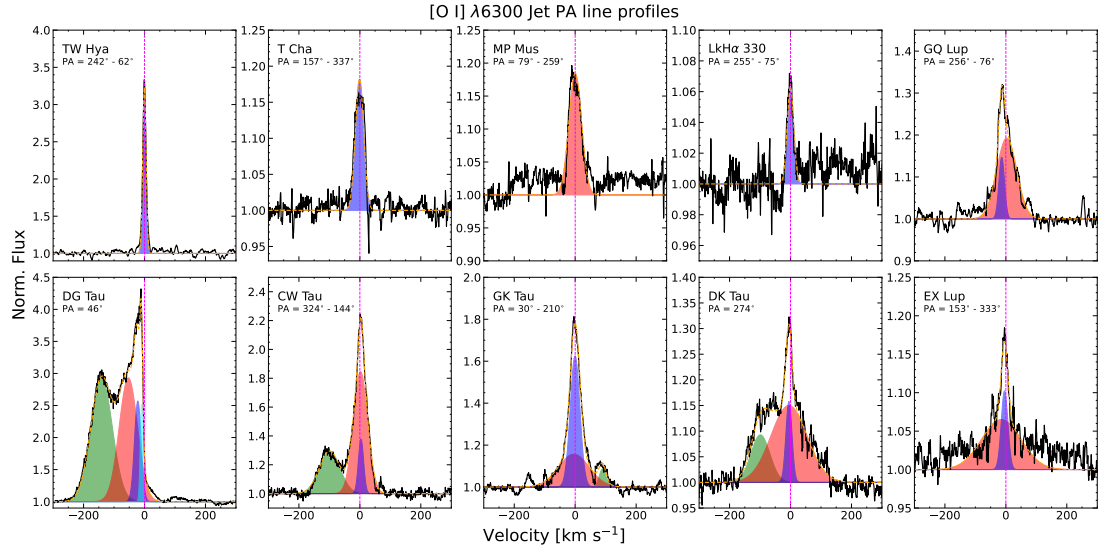


FIGURE 4.5: Comparison of the [O I] $\lambda 6300$ line along the jet axis for each target in the stellocentric frame. The spectra have been normalised to the continuum level. The dashed magenta line corresponds to 0 km s^{-1} . Green = HVC, Red = LVC-BC, Blue = LVC-NC. In the case of DG Tau, two LVC-NCs were detected in the data from 2021. As such the second LVC-NC has been coloured cyan.

[S II] $\lambda 6731$ line is detected in the spectrum of GQ Lup. For EX Lup, the [S II] $\lambda 6731$ line is only marginally detected and no LVC is present. [O I] $\lambda 6300$ line profiles of both GQ Lup and EX Lup are reproduced using a superposition of a LVC-BC and LVC-NC only. It is observed that only stars that show both a clear [O I] $\lambda 6300$ and [S II] $\lambda 6731$ line profile simultaneously, seem to have a HVC sub-component in their [O I] $\lambda 6300$ line profiles. The HVC is a long known tracer of high velocity jets in CTTs (Frank et al., 2014; Ray et al., 2007) so, observing both the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ lines in these spectra seems to indicate that different layers of an extended jet are being traced for these stars.

TABLE 4.3: Results of the kinematic fitting procedure for the [O I] $\lambda 6300$ line profiles shown in Fig. 4.5. In all cases, the uncertainty calculated using the method of Nisini et al. (2024) was found to be smaller than the uncertainty on the radial velocity for each target. As such, the uncertainty for each value of V_c and FWHM is taken to be the smallest uncertainty on the radial velocity given in Table 2.2.

Target	Comp. Type	V_c [km s ⁻¹]	FWHM [km s ⁻¹]
TW Hya	LVC-NC	0 ± 0.3	13 ± 0.3
T Cha	LVC-NC	1 ± 0.5	32 ± 0.5
MP Mus	LVC-BC	0 ± 0.4	48 ± 0.4
LkH α 330	LVC-NC	0 ± 0.7	25 ± 0.7
GQ Lup	LVC-NC	-14 ± 0.3	22 ± 0.3
	LVC-BC	3 ± 0.3	77 ± 0.3
DG Tau	LVC-NC	-10 ± 0.5	10 ± 0.5
	LVC-NC	-22 ± 0.5	26 ± 0.5
	LVC-BC	-53 ± 0.5	71 ± 0.5
	HVC	-141 ± 0.5	83 ± 0.5
CW Tau	LVC-NC	4 ± 0.4	20 ± 0.4
	LVC-BC	2 ± 0.4	55 ± 0.4
	HVC	-99 ± 0.4	80 ± 0.4
GK Tau	LVC-NC	0 ± 0.3	37 ± 0.3
	LVC-BC	-4 ± 0.3	119 ± 0.3
	HVC	90 ± 0.3	26 ± 0.3
DK Tau	LVC-NC	-5 ± 0.4	23 ± 0.4
	LVC-BC	-5 ± 0.4	139 ± 0.4
	HVC	-99 ± 0.4	73 ± 0.4
EX Lup	LVC-NC	-4 ± 0.4	26 ± 0.4
	LVC-BC	-15 ± 0.4	152 ± 0.4

4.3.2 Spectro-astrometric analysis

This section presents the spectro-astrometric analysis (position spectra and Δ FWHM spectra) of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs present in this sample. For posterity the spectro-astrometric analysis of the HVCs present in the sample are also shown but, the discussion of the spectro-astrometric analysis of these HVCs is beyond the scope

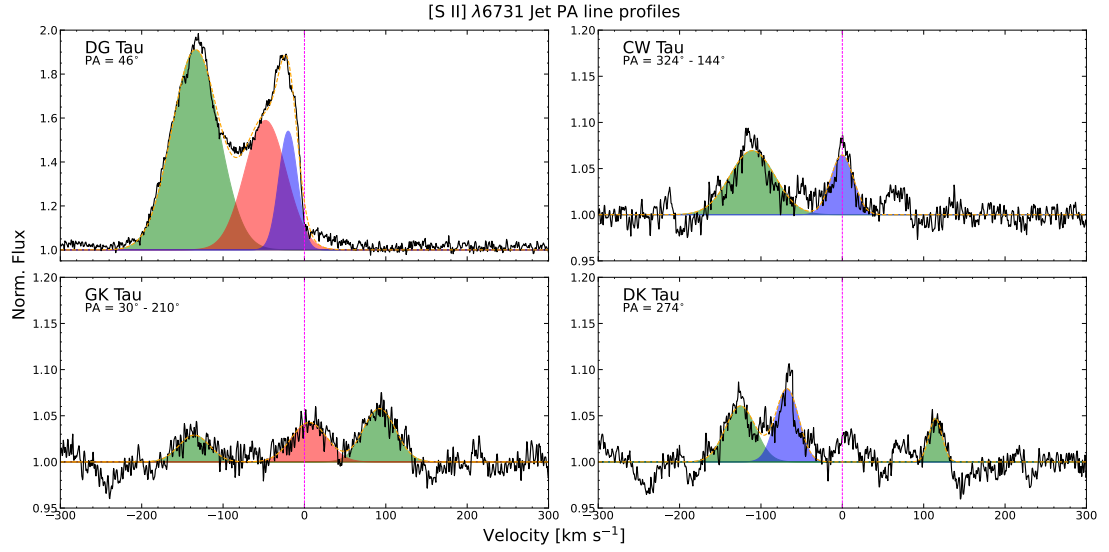


FIGURE 4.6: Comparison of the [S II] $\lambda 6731$ lines along the jet axis for each target showing this emission line in the stellocentric frame. The spectra have been normalised to the continuum level. The dashed magenta line corresponds to 0 km s^{-1} . Green = HVC, Red = LVC-BC, Blue = LVC-NC.

of the study presented in this chapter. In addition to this, given the argument presented in Section 4.2.3, the spectro-astrometric analysis for the PAs along the jet **only** are shown. The disc PAs were checked for every CTTs in the sample as shown above for TW Hya and no real signal was detected for the LVC along the disc PAs. Lastly, it is noted that for some sources (e.g. DG Tau) the anti-parallel subtraction procedure was not performed as the corresponding anti-parallel PA spectrum was not taken during the observation runs.

Figures 4.7 to 4.9 presents the analysis of all the SC sources in this sample as well as the analysis for GQ Lup and EX Lup in [O I] $\lambda 6300$. For all of these sources it was determined that the offset measured in the position spectra (middle panels) does not "flip" between anti-parallel jet PAs (as shown for TW Hya above). Given this, the value of σ_{avg} presented in the position spectra (middle panels in these figures) was used as a lower limit estimate of the vertical extent of the [O I] $\lambda 6300$ emission above the disc mid-plane in for the LVC sub-components. Similarly, the $\Delta FWHM$ spectra (bottom panels) do not present a clear signal for these stars in the [O I] $\lambda 6300$ emission line.

TABLE 4.4: Results of the kinematic fitting procedure for the [S II] $\lambda 6731$ line profiles shown in Fig. 4.6. In all cases, the uncertainty calculated using the method of Nisini et al. (2024) was found to be smaller than the uncertainty on the radial velocity for each target. As such, the uncertainty for each value of V_c and FWHM is taken to be the smallest uncertainty on the radial velocity given in Table 2.2.

Target	Comp. Type	V_c [km s ⁻¹]	FWHM [km s ⁻¹]
DG Tau	LVC-NC	-20 ± 0.5	27 ± 0.5
	LVC-BC	-48 ± 0.5	60 ± 0.5
	HVC	-134 ± 0.5	68 ± 0.5
CW Tau	LVC-NC	0 ± 0.4	33 ± 0.4
	HVC	-111 ± 0.4	65 ± 0.4
GK Tau	LVC-BC	59 ± 0.3	51 ± 0.3
	HVC	-136 ± 0.3	45 ± 0.3
	HVC	92 ± 0.3	47 ± 0.3
DK Tau	MVC*	-68 ± 0.4	34 ± 0.4
	HVC	-125 ± 0.4	44 ± 0.4
	HVC	115 ± 0.4	19 ± 0.4

* = This kinematic component has been designated as a medium velocity component (MVC) given that its peak velocity lies between the cut-off for either an LVC or HVC (see Section 1.2.5).

Figures 4.10 to 4.13 show the spectro-astrometric analysis for DG Tau, CW Tau, GK Tau and DK Tau for both the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs. For CW Tau, GK Tau and DK Tau the expected "flip" between the anti-parallel PAs doesn't occur for the LVC sub-components in [O I] $\lambda 6300$. However, for DG Tau, a clear offset is observed in all of its LVC sub-components in agreement with Chou et al. (2025). Considering the lack of anti-parallel PA subtraction (due to the lack of a spectrum taken at a PA of 226°), it was assumed that these offsets measured in the position spectra are real given that the offsets measured are well above $3\sigma_{avg}$ in both the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs. For the [S II] $\lambda 6731$ line of CW Tau and GK Tau, no real offset is observed for the LVC sub-components. This may be attributed to the continuum subtraction. The flux of the LVC sub-component is diminished following the continuum subtraction on the 2D spectra. This decrease in flux causes a drop in the S/N ratio in LVC region of the line. Lower S/N makes the spatial profiles in LVC region of the line less Gaussian

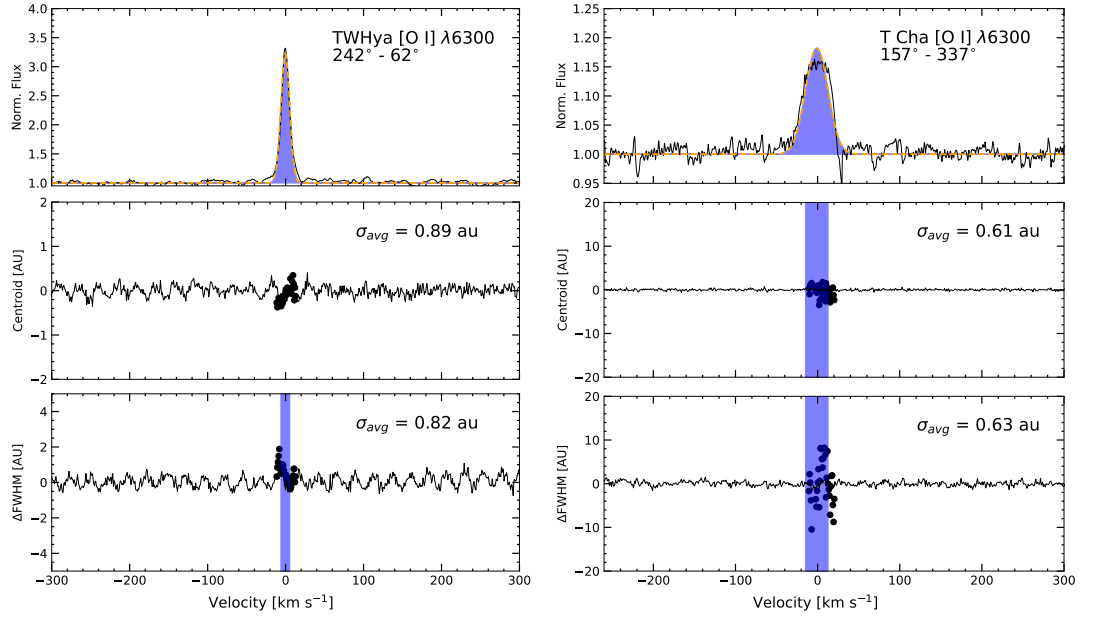


FIGURE 4.7: Spectroastrometric analysis of TW Hya (left) and T Cha (right) for [O I] $\lambda 6300$ along the jet axis. Top panels: Continuum normalised line profile. Middle Panels: Position spectra. Bottom Panels: Δ FWHM spectra. The solid lines in the middle and bottom panels are prior to continuum subtraction. The black points are after continuum subtraction.

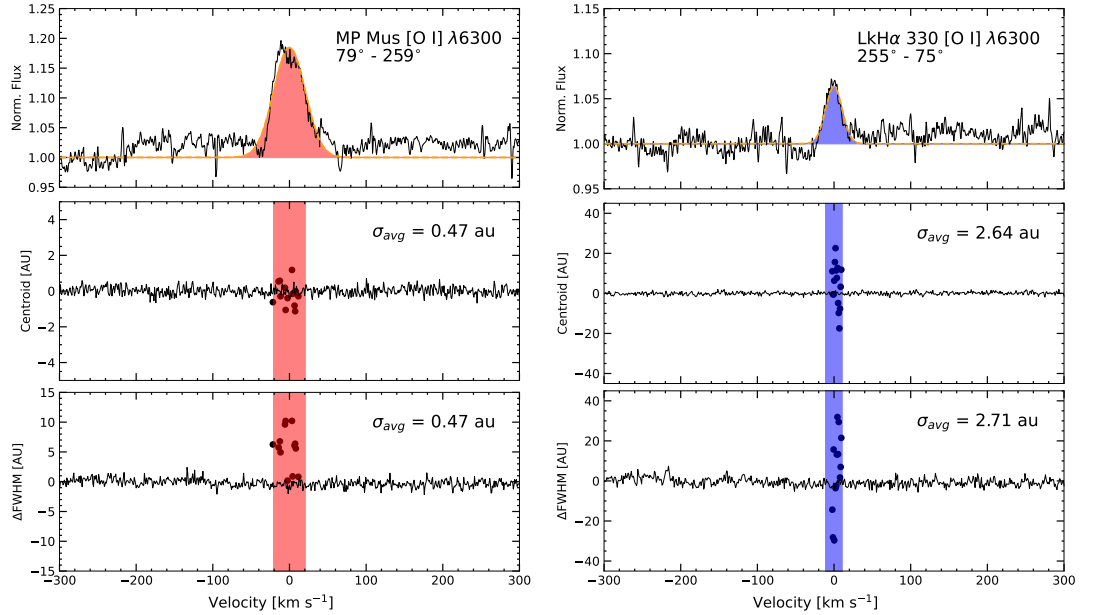


FIGURE 4.8: Spectroastrometric analysis of MP Mus (left) and LkH α 330 for [O I] $\lambda 6300$ along the jet axis.

shaped, and therefore the Gaussian fit to these spatial profiles are unreliable. Hence, the centroid and Δ FWHM values cannot be extracted from the LVC of [S II] $\lambda 6731$ line

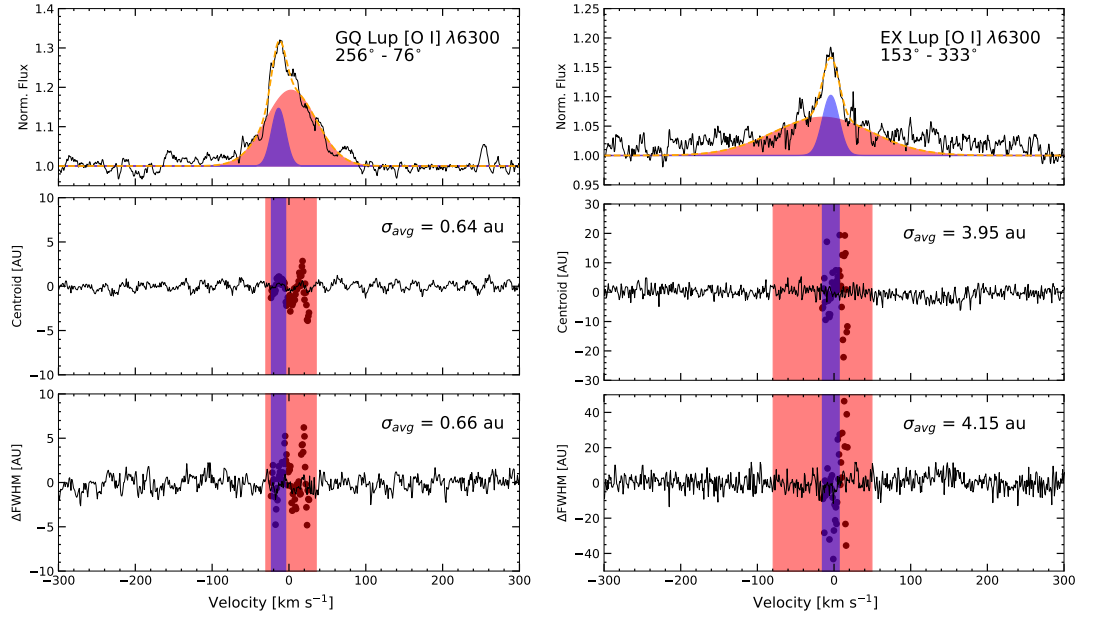


FIGURE 4.9: Spectroastrometric analysis of GQ Lup (left) and EX Lup (right) for [O I] $\lambda 6300$ along the jet axis.

for CW Tau and GK Tau following the continuum subtraction. Moving to DK Tau, there appears to be a peak offset of ~ 80 AU in the [S II] $\lambda 6731$ position spectra of this star but, the signal seems to be blended between the HVC (marked as a green shaded region) and the MVC (marked as a blue shaded region) in the position spectrum. The $\Delta FWHM$ spectrum of DK Tau in [S II] $\lambda 6731$ appears very scattered. Given this, minimal information regarding the intrinsic size of the emission region can be obtained from this analysis.

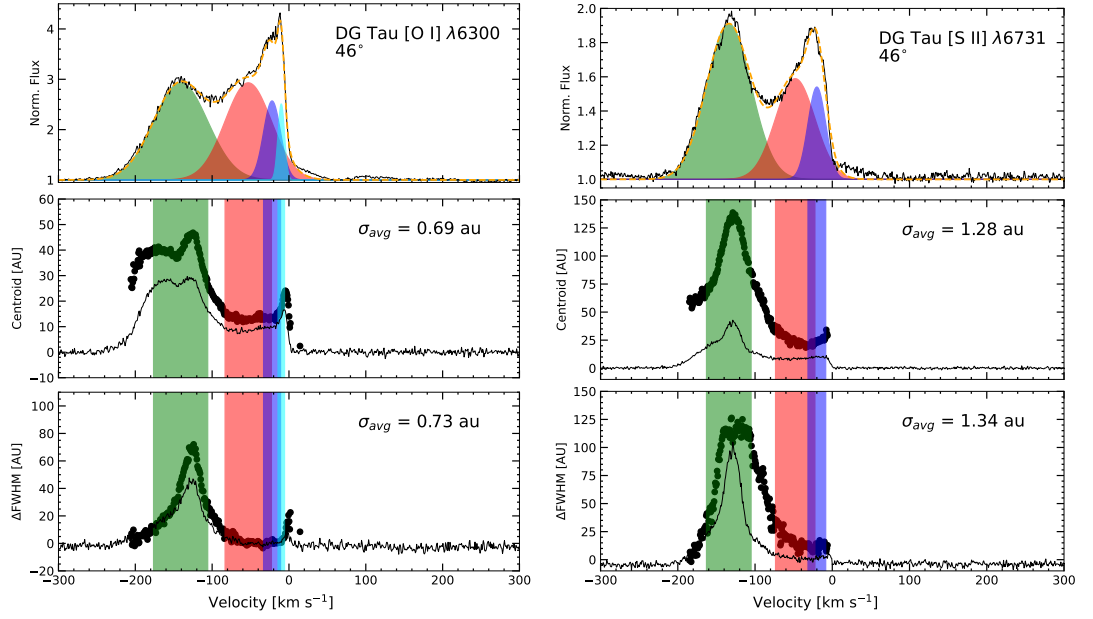


FIGURE 4.10: Spectroastrometric analysis of DG Tau for the [O I] $\lambda 6300$ (left) and [S II] $\lambda 6731$ (right) along the jet axis.

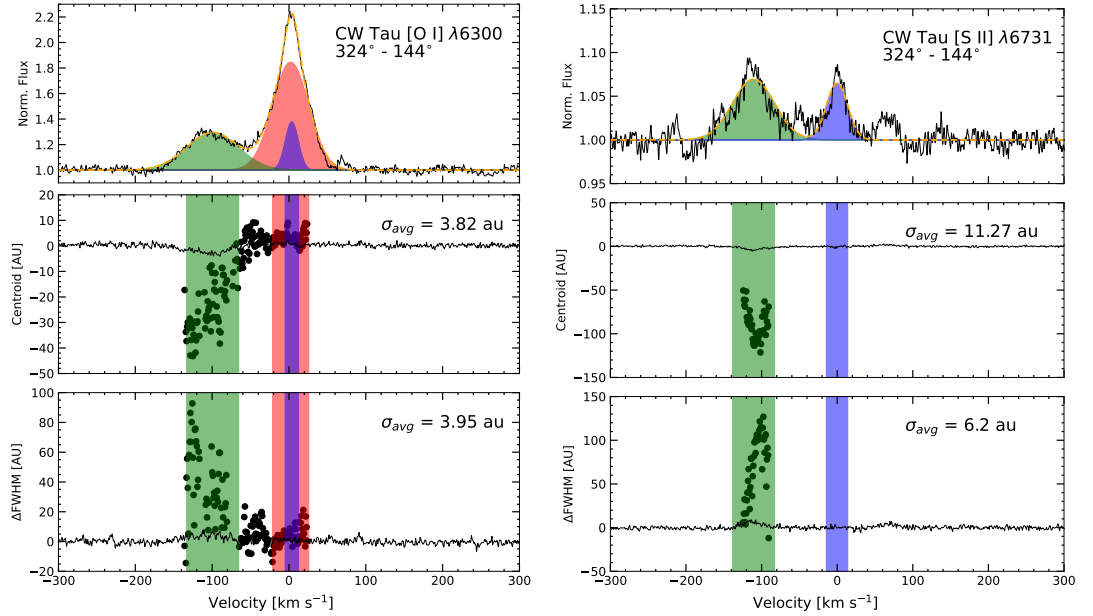


FIGURE 4.11: Spectroastrometric analysis of CW Tau for the [O I] $\lambda 6300$ (left) and [S II] $\lambda 6731$ (right) along the jet axis.

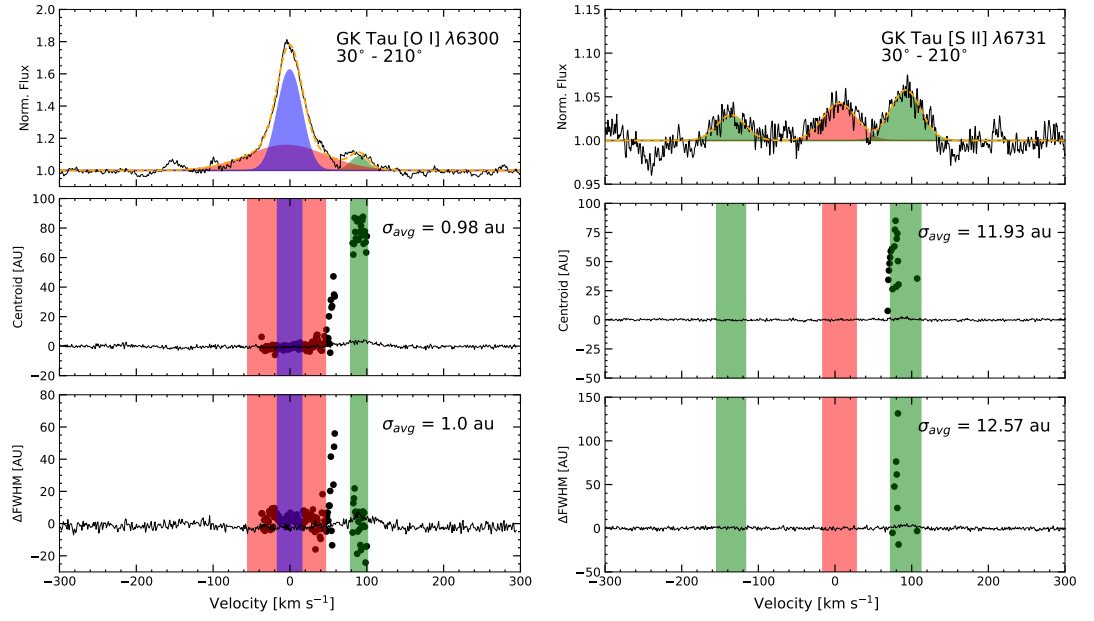


FIGURE 4.12: Spectroastrometric analysis of GK Tau for the [O I] $\lambda 6300$ line (left) and the [S II] $\lambda 6731$ line (right) along the jet axis.

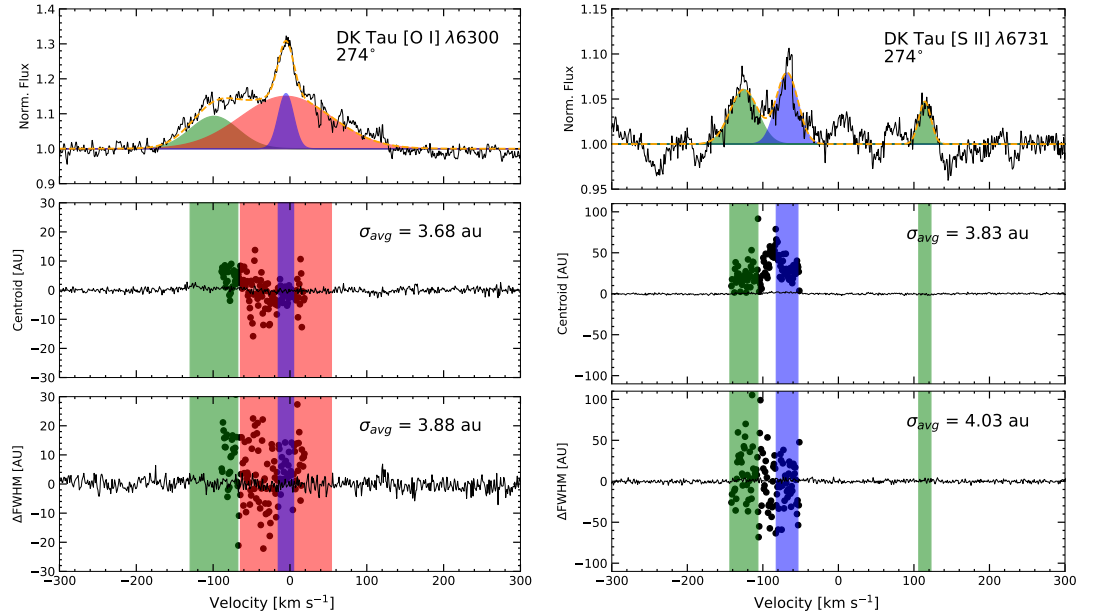


FIGURE 4.13: Spectroastrometric analysis of DK Tau for the [O I] $\lambda 6300$ line (left) and the [S II] $\lambda 6731$ line (right) along the jet axis.

The blue shaded region for the [S II] $\lambda 6731$ corresponds to a MVC.

4.3.3 Outflow efficiency, mass accretion rates and mass outflow rates for the LVC

Using the combined information gleaned from both kinematic fitting and spectro-astrometric analysis, a test of a key assumption of disc wind driven accretion models may be preformed. Namely, that the mass outflow rate from the low velocity wind is comparable to the stellar mass accretion rate. This assumption arises from recent simulations (e.g. Bai, 2016) and the realisation that a low velocity MHD disc wind could play a dominant role in the angular momentum transport of the outer disc regions (see Pascucci et al., 2023, and references therein). Given this, it is expected that the mass lost due to an MHD disc wind is inherently linked to angular momentum transport in the disc. Alongside this, the work of Hartmann, Herczeg, and Calvet (2016) has also shown that angular momentum removal in these outer disc regions is a requirement for mass to be transported towards the inner disc regions and ultimately accreted by the YSO. Therefore, given a typical disc lifetime of a few million years, it is expected that the rate at which the wind extracts mass (and therefore disc angular momentum) is comparable to the rate at which the YSO accretes mass.

Outflow efficiency is defined as the ratio of the mass outflow rate to the stellar mass accretion rate as shown in Equation 4.1. This is a key parameter that must be derived in order to test current disc wind models (Bai, 2016, 2017). In order for the current disc winds models to hold, an outflow efficiency approaching unity is required for the low velocity flow traced by the LVC.

$$\eta = \frac{\dot{M}_{out}}{\dot{M}_{acc}} \quad (4.1)$$

Previous studies of high velocity jets (Cabrit et al., 2010; Dougados et al., 2000; Fang et al., 2018) have derived an outflow efficiency in the order $\sim 10\%$ for these high velocity ejecta, however, the efficiency of the low velocity flows remains more uncertain. The works of Whelan et al. (2021) and Birney et al. (2024b) have recently found that for the CTTS RU Lupi, the efficiency of the LVC-NC is considerably lower than that of the

high velocity jet. It is shown to be a fraction of a percent. This is in contradiction to the assumptions of current disc wind models. However, it is unclear if the findings of these recent studies hold in general for a larger sample of CTTSs or if RU Lupi is unique. Another study by Salyk et al. (2014) concerning the CTTS AS205 N has estimated an efficiency of ≈ 1 to 10% using ALMA, albeit with large uncertainty. This section aims to expand on these recent works by deriving a stellar mass accretion, mass outflow and outflow efficiency values for each CTTS in each LVC sub-component presented above.

Mass accretion rates

Of the ten CTTSs in the sample, seven had flux calibrated spectra. The stellar mass accretion rate for the flux calibrated targets in the sample was derived using the method described in Alcalá et al., 2017. For the three remaining non flux calibrated targets (TW Hya, T Cha and MP Mus), the methods presented in Whelan et al. (2009) and Pascucci et al. (2020) were used to convert the Equivalent Width (EW) of the [O I] $\lambda 6300$ line into a line luminosity which can later be converted to an accretion luminosity and a subsequent stellar mass accretion rate. This subsection now provides a short description of each of these methods followed by the presentation of the derived stellar mass accretion rates.

For each of the flux calibrated target spectra, observations taken on the same night with the same PA and comparable exposure times were median combined using the *imcombine* task in IRAF. The integrated line flux was then measured in IRAF using the *splot* task for a handful of permitted lines that are known accretion tracers (i.e. $H\alpha$, $He\lambda 5875$, Na D lines).

To correct the measured permitted line flux for line of sight extinction the *Python* package *PyNeb* (Luridiana, Morisset, and Shaw, 2015) was used to compute a multiplicative extinction correction factor for a given input flux and A_v value taken from the literature (see Table 4.1). Following extinction correction, the corrected line flux was

converted to a line luminosity following Equation 4.2 where f_{line} is the extinction corrected line flux in $\text{ergs s}^{-1} \text{cm}^2$ and d is the distance to the target star in centimetres.

$$L_{line} = 4\pi d^2 \times f_{line} \quad (4.2)$$

The line luminosity is subsequently converted into an accretion luminosity using the relation shown by Equation 4.3, where L_{acc} is the accretion luminosity and the values of a and b are coefficients determined by Alcalá et al. (2017) via linear fits to $\log L_{acc}$ vs $\log L_{line}$ from a large sample of low mass YSOs.

$$\log(L_{acc}/L_{\odot}) = a \times \log(L_{line}/L_{\odot}) + b \quad (4.3)$$

Finally, the accretion luminosity is converted to a stellar mass accretion rate in units of $M_{\odot}/\text{yr}$ following Equation 4.4, where G is the universal gravitational constant, R_{in} is the inner disc radius and $M_{\star}$ and $R_{\star}$ are the assumed mass and radius of the star. For consistency with previous works (e.g. Alcalá et al., 2017; Fang et al., 2018), R_{in} is assumed to be $5 R_{\star}$.

$$\dot{M}_{acc} = \frac{L_{acc} R_{\star}}{G M_{\star} (1 - \frac{R_{\star}}{R_{in}})} \approx 1.25 \frac{L_{acc} R_{\star}}{G M_{\star}} \quad (4.4)$$

Table 4.5 presents the measured line fluxes, line luminosities, accretion luminosities and mass accretion rates derived for each of the seven flux calibrated CTTSs in each permitted line accretion tracer using the method of Alcalá et al. (2017). An average stellar mass accretion rate for each star was found by averaging the mass accretion rates value found from each permitted line. These average mass accretion rates for each CTTS were then compared to an existing value of stellar mass accretion rate in the literature for posterity. These averaged stellar mass accretion rates presented in Table 4.6 will be used in further calculations later for the outflow efficiency rates.

As mentioned previously, three of the ten targets selected for this study (i.e. TW Hya, T Cha and MP Mus) did not have any flux calibrated spectra available in either of the chosen observing programmes. Given this, an alternative method was used to

TABLE 4.5: Stellar mass accretion rates for the sample of CTTs using the method of Alcalá et al. (2017).

Target	Line	a,b	Integrated Flux [erg s ⁻¹ cm ⁻²]	Correction Factor	Corrected Flux [erg s ⁻¹ cm ⁻²]	Log(L _{line}) [L _☉]	Log(L _{acc}) [L _☉]	Log($\dot{M}_{acc}$)
LkHα 330	He I $\lambda 5875$	1.15,3.67	1.09×10^{-14}	4.66	5.08×10^{-14}	-3.79	-0.69	-8.26
	Na I $\lambda 5890$	1.01,3.14	1.31×10^{-14}	4.64	6.07×10^{-14}	-3.72	-0.61	-8.18
	Na I $\lambda 5896$	1.01,3.33	1.61×10^{-14}	4.63	7.45×10^{-14}	-3.63	-0.33	-7.90
	H α	1.13,1.74	5.84×10^{-13}	3.88	2.27×10^{-12}	-2.14	-0.68	-8.25
GQ Lup	He I $\lambda 5875$	1.15,3.67	4.19×10^{-14}	1.53	6.42×10^{-14}	-4.32	-1.30	-8.39
	Na I $\lambda 5890$	1.01,3.14	1.38×10^{-14}	1.53	2.11×10^{-14}	-4.80	-1.71	-8.81
	Na I $\lambda 5896$	1.01,3.33	2.18×10^{-14}	1.53	3.34×10^{-14}	-4.61	-1.32	-8.42
	H α	1.13,1.74	3.67×10^{-12}	1.46	5.35×10^{-12}	-2.40	-0.97	-8.07
DG Tau	He I $\lambda 5875$	1.15,3.67	9.48×10^{-14}	3.93	3.72×10^{-13}	-3.74	-0.63	-7.67
	Na I $\lambda 5890$	1.01,3.14	2.34×10^{-14}	3.91	9.15×10^{-14}	-4.35	-1.25	-8.29
	Na I $\lambda 5896$	1.01,3.33	4.98×10^{-14}	3.90	1.94×10^{-13}	-4.02	-0.73	-7.77
	H α	1.13,1.74	3.18×10^{-12}	3.36	1.06×10^{-11}	-2.28	-0.84	-7.88
CW Tau	He I $\lambda 5875$	1.15,3.67	6.16×10^{-14}	10.49	6.46×10^{-13}	-3.46	-0.31	-7.56
	Na I $\lambda 5890$	1.01,3.14	4.69×10^{-14}	10.42	4.89×10^{-13}	-3.58	-0.47	-7.73
	Na I $\lambda 5896$	1.01,3.33	2.27×10^{-14}	10.39	2.36×10^{-13}	-3.89	-0.60	-7.86
	H α	1.13,1.74	4.09×10^{-12}	7.94	3.25×10^{-11}	-1.76	-0.24	-7.50
GK Tau	He I $\lambda 5875$	1.15,3.67	2.53×10^{-14}	2.23	5.65×10^{-14}	-4.53	-1.54	-8.52
	Na I $\lambda 5890$	1.01,3.14	2.47×10^{-14}	2.23	5.50×10^{-14}	-4.54	-1.45	-8.43
	Na I $\lambda 5896$	1.01,3.33	1.21×10^{-14}	2.23	4.92×10^{-14}	-4.59	-1.31	-8.29
	H α	1.13,1.74	1.13×10^{-12}	2.03	2.29×10^{-12}	-2.92	-1.56	-8.55
DK Tau	He I $\lambda 5875$	1.15,3.67	1.33×10^{-14}	3.04	4.04×10^{-14}	-4.66	-1.69	-8.65
	Na I $\lambda 5890$	1.01,3.14	8.29×10^{-15}	3.03	2.51×10^{-14}	-4.86	-1.77	-8.74
	Na I $\lambda 5896$	1.01,3.33	9.47×10^{-15}	3.02	2.86×10^{-14}	-4.81	-1.53	-8.49
	H α	1.13,1.74	9.09×10^{-13}	2.66	2.42×10^{-12}	-2.88	-1.51	-8.48
EX Lup	He I $\lambda 5875$	1.15,3.67	1.72×10^{-14}	2.56	4.40×10^{-14}	-4.48	-1.48	-8.38
	Na I $\lambda 5890$	1.01,3.14	6.62×10^{-15}	2.55	1.69×10^{-14}	-4.90	-1.81	-8.70
	Na I $\lambda 5896$	1.01,3.33	3.93×10^{-15}	2.55	1.00×10^{-14}	-5.13	-1.85	-8.74
	H α	1.13,1.74	6.52×10^{-13}	2.29	1.49×10^{-12}	-2.95	-1.60	-8.49

derive the stellar mass accretion rates for these CTTs. These methods, described in Whelan et al. (2009) and Pascucci et al. (2020) use the relations shown by Equations 4.6 and 4.7 respectively to calculate an [O I] $\lambda 6300$ line luminosity given the EW of the [O I] $\lambda 6300$ line, the distance to the target, the R or V band magnitude for the target and an estimate for the extinction towards the target in each band. The EW is defined as shown in Equation 4.5, where I_{cont} is the continuum intensity and I_{λ} is the line intensity at a wavelength λ .

$$EW = \int \frac{I_{cont} - I_{\lambda}}{I_{cont}} d\lambda \quad (4.5)$$

The EW of the [O I] $\lambda 6300$ line of each of the ten targets were measured in IRAF by fitting a Gaussians to each line sub-component using the *splot* routine. From this, assuming that the spectra are wavelength calibrated, IRAF then estimates the EW of the

TABLE 4.6: Averaged stellar mass accretion rates found using the method of Alcalá et al. (2017) for the flux calibrated target CTTs in this sample compared to existing values in literature.

Target	Lit. $\text{Log}(\dot{M}_{\text{acc}})$ [$M_{\odot}/\text{yr}$]	Avg. Meas. $\text{Log}(\dot{M}_{\text{acc}})$ [$M_{\odot}/\text{yr}$]	Ref.
TW Hya	-8.81	-	HH08
T Cha	-8.40	-	S09
MP Mus	-8.30	-	P07
LkH α 330	-8.80	-8.15	Sa09
GQ Lup	-9.50	-8.42	A20
DG Tau	-8.39	-7.90	HH08
CW Tau	-7.47	-7.66	G22
GK Tau	-8.19	-8.45	Ha98
DK Tau	-8.33	-8.59	G22
EX Lup	-7.44	-8.58	A17

A17 = Alcalá et al. (2017); A20 = Alcalá et al. (2020); G22 = Gangi et al. (2022); Ha98 = Hartmann et al. (1998); HH08 = Herczeg and Hillenbrand (2008); P07 = Pascucci et al. (2007); S09 = Schisano et al. (2009); Sa09 = Salyk et al. (2009); Sa13 = Salyk et al. (2013).

[O I] $\lambda 6300$ line for each star using Equation 4.5.

The R and V band magnitudes for each target was taken directly from measurements available within the SIMBAD database (Wenger et al., 2000). If R and V band magnitudes were not available for a specific star within SIMBAD, the American Association of Variable Star Observers (AAVSO) database (available at <https://www.aavso.org/>) was used as a second option. If either database did not contain a specific R or V band magnitude for a given target then the literature was searched in order to acquire a magnitude for that band/star. For LkH α 300, no R band magnitude was available in either database or from the literature search so, an [O I] $\lambda 6300$ line luminosity (using the [O I] $\lambda 6300$ EW) could not be calculated for this stars using the R band. These magnitude values were corrected for line of sight extinction using *PyNeb* similarly to the targets with available flux calibrated spectra. For the V band magnitudes, visual extinction values, A_v (see Table 4.1), were used to calculate a multiplicative factor to correct for line of sight extinction. In the R band, the A_v values obtained from the literature were first converted to extinction values in the R band (A_R) using the relation $A_R \sim 0.79 \times A_v$ (Fitzpatrick (1999)); assuming a total to selective extinction ratio

of $R \sim 3.1$; which is the typical value used for the diffuse ISM). Once converted, the A_R value was used in *PyNeb* to calculate a multiplicative factor to correct the R band magnitudes for the line of sight extinction.

Using the measured values of EW for the [O I] $\lambda 6300$ line, the extinction corrected R and V band magnitudes and the assumed distance to each star (see Table 4.1); Equations 4.6 and 4.7 were used to calculate an [O I] $\lambda 6300$ line luminosity for each star. The units for each of the quantities in these Equations are specified in brackets beside each quantity. Vega mags are used as the unit for the R and V band magnitudes and there respective extinction values. Table 4.7 presents the line luminosity values for the [O I] $\lambda 6300$ line of each star calculated using the EW of the [O I] $\lambda 6300$ line.

TABLE 4.7: [O I] $\lambda 6300$ line luminosity values for both the flux and non flux calibrated targets in the sample using the methods outlined in Whelan et al. (2009) and Pascucci et al. (2020).

Target	Comp. Type	Equiv. Width [Å]	Rmag [mag]	Vmag [mag]	A_R [mag]	$\text{Log}(L_{6300, \text{Whelan}})$ [$L_\odot$]	$\text{Log}(L_{6300, \text{Pascucci}})$ [$L_\odot$]	Refs. ^a
TW Hya	LVC-NC	-0.327	10.188 ± 0.008 ^b	10.887 ± 0.005	0.0	-5.176	-5.296	Ro05; AAVSO
T Cha	LVC-NC	-0.135	10.533 ± 0.006	11.421 ± 0.006	1.3	-4.696	-4.748	AAVSO; AAVSO
MP Mus	LVC-NC	-0.173	~ 10 ^b	10.393 ± 0.068	0.6	-4.733	-4.671	Av18; SIMBAD
LkH α 330	LVC-NC	-0.053	$\sim b, c$	10.970	1.4	–	-3.953	–; SIMBAD
GQ Lup	LVC-NC	-0.224	11.150 ± 0.120	11.660 ± 0.400	0.4	-4.749	-4.752	SIMBAD; SIMBAD
	LVC-BC	-0.304	11.150 ± 0.120	11.660 ± 0.400	0.4	-4.617	-4.619	SIMBAD; SIMBAD
DG Tau	LVC-NC1	-1.206	12.280 ^d	13.035 ± 0.009	1.3	-4.303	-4.311	SIMBAD; AAVSO
	LVC-NC2	-1.713	12.280 ^d	13.035 ± 0.009	1.3	-4.150	-4.158	SIMBAD; AAVSO
	LVC-BC	-5.066	12.280 ^d	13.035 ± 0.009	1.3	-3.679	-3.687	SIMBAD; AAVSO
	HVC	-4.602	12.280 ^d	13.035 ± 0.009	1.3	-3.721	-3.729	SIMBAD; AAVSO
CW Tau	LVC-NC	-1.123	11.585 ± 0.004	13.530 ± 0.001	2.2	-3.634	-4.017	AAVSO; AAVSO
	LVC-BC	-1.382	11.585 ± 0.004	13.530 ± 0.001	2.2	-3.544	-3.927	AAVSO; AAVSO
	HVC	-0.507	11.585 ± 0.004	13.530 ± 0.001	2.2	-3.979	-4.362	AAVSO; AAVSO
GK Tau	LVC-NC	-1.283	12.020 ± 0.140	13.000 ^d	0.7	-4.367	-4.523	SIMBAD; SIMBAD
	LVC-BC	-1.331	12.020 ± 0.140	13.000 ^d	0.7	-4.351	-4.507	SIMBAD; SIMBAD
	HVC	-0.350	12.020 ± 0.140	13.000 ^d	0.7	-4.931	-5.087	SIMBAD; SIMBAD
DK Tau	LVC-NC	-0.090	12.080 ± 0.480	13.050 ± 0.580	1.0	-5.399	-5.518	SIMBAD; SIMBAD
	LVC-BC	-0.255	12.080 ± 0.480	13.050 ± 0.580	1.0	-4.947	-5.066	SIMBAD; SIMBAD
	HVC	-0.218	12.080 ± 0.480	13.050 ± 0.580	1.0	-5.015	-5.134	SIMBAD; SIMBAD
EX Lup	LVC-NC	-0.115	12.684 ± 0.054	13.510 ± 0.090	0.9	-5.459	-5.538	AAVSO; AAVSO
	LVC-BC	-0.283	12.684 ± 0.054	13.510 ± 0.090	0.9	-5.068	-5.147	AAVSO; AAVSO

References: Av18 = Avenhaus et al. (2018); Ro05 = Roberge, Weinberger, and Malumuth (2005).

^a = The references are for the R and V band magnitudes are displayed here respectively delimited by a semicolon (;). SIMBAD = The SIMBAD database (Wenger et al., 2000). AAVSO = The AAVSO database (available at <https://www.aavso.org/>).

^b = No R band magnitude available in either the SIMBAD or AAVSO databases, so another reference was used.

^c = Several sources (including an extensive literature search) were checked (i.e. SIMBAD, AAVSO, and the catalogue of Herbig and Bell (1995); which quotes a V band magnitude of 11.2) were searched for an R band magnitude. From these, an R band magnitude could not be determined for this source.

^d = The magnitude value provided within the SIMBAD database doesn't have a quoted uncertainty and no equivalent value was available within the AAVSO database.

$$L_{6300} (L_\odot) = 6.71 \times 10^{-5} \times D^2(pc) \times EW(\text{\AA}) \times 10^{-0.4(R-A_R)} \quad (4.6)$$

$$L_{6300} \text{ (erg s}^{-1}\text{)} = 3.1 \times 10^{-9} \times 4\pi D^2(\text{cm}) \times EW(\text{\AA}) \times 10^{-0.4(V - A_v)} \quad (4.7)$$

The [O I] λ 6300 line luminosities calculated in Table 4.7 (i.e using the [O I] λ 6300 EW) they were then converted into accretion luminosities and subsequently mass accretion rates using Equations 4.3 and 4.4 respectively. The a and b coefficients adapted for Equation 4.3 for [O I] λ 6300 are taken from Nisini et al. (2018) and are $a = 0.59 \pm 0.04$ and $b = -4.13 \pm 0.11$ for the [O I] λ 6300 LVC-NC and LVC-BC and $a = 0.75 \pm 0.08$ and $b = -4.03 \pm 0.17$ for the [O I] λ 6300 HVC respectively. These coefficients were chosen over the [O I] λ 6300 a, b coefficients provided in Alcalá et al. (2017) as the [O I] λ 6300 coefficients are updated compared to the values provided by Alcalá et al. (2017) and as such, these newer coefficients both account for the fact that the [O I] λ 6300 line is divided into sub-components and also that the [O I] λ 6300 line is an outflow tracer. The accretion luminosities and mass accretion rates calculated for each target using the EW methods of Pascucci et al. (2020) and Whelan et al. (2009) are presented in Table 4.8 and are compared to previously derived literature values.

Lastly, it is noted here that the mass accretion rates derived for TW Hya, T Cha and MP Mus are not presented in Table 4.6. This is because these values are not derived via the line flux method and as such, are not averaged over multiple emission lines. Comparing the methods of Whelan et al. (2009) and Pascucci et al. (2020) in using the EW to derive the stellar mass accretion rate for each star, it was observed that they are in agreement; with < 0.4 dex difference between any two measurements for the same star. In contrast, the mass accretion rate estimates derived from the two separate methods (i.e. the line flux method and the EW method) to each other, it was found that in general the EW method gives higher mass accretion rate estimate for each star compared to the line flux method by up to a few orders of magnitude. In general this implies that the EW method overestimates the stellar mass accretion rate when comparing these estimates to literature values. Given this, the mass accretion rate estimates for TW Hya, T Cha and MP Mus using the methods of Whelan et al. (2009) and Pascucci et al. (2020) will **not** be used in further calculations of outflow efficiency

TABLE 4.8: Stellar mass accretion rate estimates for the both the flux and non flux calibrated sources in the sample using the [O I] $\lambda 6300$ line luminosities derived in Table 4.7 compared to the literature.

Target	$\text{Log}(L_{6300})$ [$L_{\odot}$]	$\text{Log}(L_{acc})$ [$L_{\odot}$]	Meas. $\text{Log}(\dot{M}_{acc})^a$ [$M_{\odot}/\text{yr}$]	Lit. $\text{Log}(\dot{M}_{acc})^b$ [$M_{\odot}/\text{yr}$]
TW Hya	-5.176/ -5.296	-1.772/ 1.975	-7.878/ -8.081	-8.81
T Cha	-4.696/ -4.748	-0.959/ -1.048	-7.279/ -7.368	-8.40
MP Mus	-4.733/ -4.671	-1.021/ -0.917	-7.354/ -7.250	-8.30
LkH α 330	-/ -3.953	-/ 0.301	-/ -5.805	-8.80
GQ Lup: LVC-NC	-4.749/ -4.752	-1.050/ -1.054	-7.143/ -7.147	-9.50
GQ Lup: LVC-BC	-4.617/ -4.619	-0.825/ -0.829	-6.918/ -6.922	-9.50
DG Tau: LVC-NC1	-4.303/ -4.311	-0.293/ -0.306	-6.328/ -6.341	-8.39
DG Tau: LVC-NC2	-4.150/ -4.158	-0.034/ -0.048	-6.017/ -6.031	-8.39
DG Tau: LVC-BC	-3.679/ -3.687	0.764/ 0.750	-5.271/ -5.285	-8.39
DG Tau: HVC	-3.721/ -3.729	0.412/ 0.401	-5.623/ -5.634	-8.39
CW Tau: LVC-NC	-3.634/ -4.017	0.841/ 0.191	-5.417/ -6.066	-7.47
CW Tau: LVC-BC	-3.544/ -3.927	0.994/ 0.344	-5.264/ -5.913	-7.47
CW Tau: HVC	-3.979/ -4.362	0.068/ -0.443	-6.190/ -6.701	-7.47
GK Tau: LVC-NC	-4.367/ -4.523	-0.401/ -0.667	-6.384/ -6.649	-8.19
GK Tau: LVC-BC	4.351/ -4.507	-0.374/ -0.640	-6.357/ -6.622	-8.19
GK Tau: HVC	-4.931/ -5.087	-1.201/ -1.410	-7.184/ -7.393	-8.19
DK Tau: LVC-NC	-5.399/ -5.518	-2.151/ -2.353	-8.118/ -8.320	-8.33
DK Tau: LVC-BC	-4.947/ -5.066	-1.384/ -1.586	-7.351/ -7.553	-8.33
DK Tau: HVC	-5.015/ -5.134	-1.313/ -1.472	-7.280/ -7.439	-8.33
EX Lup: LVC-NC	-5.459/ -5.538	-2.253/ -2.386	-8.148/ -8.280	-7.44
EX Lup: LVC-BC	-5.068/ -5.147	-1.591/ -1.723	-7.485/ -7.617	-7.44

^a = Two values of $\dot{M}_{acc}$ are found for most target stars as two values of L_{6300} are found independently using the methods of Whelan et al. (2009) and Pascucci et al. (2020).

^b = References for the literature values are shown in Table 4.6.

in this chapter. Instead, values for the stellar mass accretion rates of TW Hya, T Cha and MP Mus taken from literature are used for the calculation of outflow efficiency. Additionally, the stellar mass accretion rate estimates for the flux calibrated stars (i.e. the measured values in Table 4.6) using the line flux method will be used in the further calculations of outflow efficiency given the clear overestimation of the stellar mass accretion rates for these stars using the EW method.

Mass outflow rates & LVC efficiency

TABLE 4.9: [O I] $\lambda 6300$ line luminosity values for each flux calibrated CTTS spectra in the sample.

Target	Comp. Type	Integrated Flux [erg s ⁻¹ cm ⁻²]	Correction Factor	Corrected Flux [erg s ⁻¹ cm ⁻²]	Log(L ₆₃₀₀) [L _⊙]
LkH α 330	LVC-NC	5.55×10^{-15}	4.15	2.30×10^{-14}	-4.14
GQ Lup	LVC-NC	4.77×10^{-15}	1.49	7.11×10^{-15}	-5.28
	LVC-BC	3.54×10^{-14}	1.49	5.28×10^{-14}	-4.41
DG Tau	LVC-NC	1.26×10^{-14}	3.54	4.46×10^{-14}	-4.66
	LVC-NC	2.64×10^{-14}	3.54	9.35×10^{-14}	-4.34
	LVC-BC	1.04×10^{-13}	3.54	3.68×10^{-13}	-3.74
	HVC	1.18×10^{-13}	3.54	4.18×10^{-13}	-3.69
CW Tau	LVC-NC	2.78×10^{-15}	8.80	2.45×10^{-14}	-4.88
	LVC-BC	1.85×10^{-14}	8.80	1.63×10^{-13}	-4.06
	HVC	1.63×10^{-14}	8.80	1.43×10^{-13}	-4.11
GK Tau	LVC-NC	1.05×10^{-14}	2.10	2.21×10^{-14}	-4.94
	LVC-BC	4.25×10^{-15}	2.10	8.93×10^{-15}	-5.33
	HVC	1.33×10^{-16}	2.10	2.79×10^{-16}	-6.84
DK Tau	LVC-NC	3.00×10^{-16}	2.80	8.40×10^{-16}	-6.34
	LVC-BC	1.28×10^{-14}	2.80	3.58×10^{-14}	-4.71
	HVC	1.07×10^{-14}	2.80	3.00×10^{-14}	-4.79
EX Lup	LVC-NC	5.93×10^{-16}	2.39	1.42×10^{-15}	-5.98
	LVC-BC	6.93×10^{-15}	2.39	1.66×10^{-14}	-4.91

In this section, the mass outflow rates for both the narrow and broad LVC sub-components are estimated. The outflow rates were derived following the method described by Fang et al. (2018) and is summarised in Equation and 4.8. Where $C(T)$ is a temperature dependant constant, V_{wind} is the velocity of the flow in units of 10s of km s⁻¹, l_{wind} is the height of the flow above the disc mid-plane in AU and L_{6300} is the [O I] $\lambda 6300$ line luminosity in solar units.

$$\dot{M}_{wind} = C(T) \times \frac{V_{wind}}{10 \text{ km s}^{-1}} \times \frac{l_{wind}}{1 \text{ au}}^{-1} \times \frac{L_{6300}}{L_{\odot}} (M_{\odot} \text{ yr}^{-1}) \quad (4.8)$$

The constant $C(T)$ varies between 2.4×10^{-4} and 2.6×10^{-5} for assumed gas temperatures of 5,000 K and 10,000 K respectively. V_{wind} is the peak velocity of each LVC sub-component obtained via the kinematic fitting presented in Fig. 4.5 and Table 4.3 deprojected to account for the disc inclination with respect to the plane of the sky. l_{wind} is the deprojected vertical height of the LVC emission above the disc mid-plane as estimated from the spectro-astrometric analysis of the [O I] $\lambda 6300$ lines shown from Fig 4.7 to 4.13. In cases where no spectro-astrometric signal is detected for a particular LVC sub-component, the value of the average spectro-astrometric uncertainty (σ_{avg}) is used as lower bound estimate for the extent of the emission traced via that sub-component. Lastly, the [O I] $\lambda 6300$ line luminosity was measured in IRAF using the *splot* task and the method of Alcalá et al. (2017) (i.e. Equation 4.3) for the flux calibrated spectra in the sample. Applying the same fit as in the kinematic fitting to each line profile yields a measured flux for each sub-component in these line profiles. This can then be corrected for extinction and converted to an accretion luminosity as described above using the linear coefficients for the [O I] $\lambda 6300$ from Alcalá et al. (2017). The [O I] $\lambda 6300$ line luminosities for each [O I] $\lambda 6300$ sub-component for each CTTS is presented in Table 4.9. The [O I] $\lambda 6300$ line luminosities for the three remaining non flux calibrated sources are shown in Table 4.7 and found using the methods described above (Pascucci et al., 2020; Whelan et al., 2009).

From these calculations, a mass outflow rate for each of the LVC sub-components can be derived. These are presented in Tables 4.10 for the broad LVC components and 4.11 for the narrow LVC components at both of the assumed gas temperatures of 5,000K and 10,000K.

The final result of this chapter is the comparison of the measured stellar mass accretion rates to the mass outflow rates in each LVC sub-component. The efficiency is defined and calculated as shown in Equation 4.1 and expressed as a percentage in the previous two tables for the broad and narrow LVC sub-components respectively. These efficiencies have been plot as a function of each sub-components FWHM in Fig. 4.14 to provide a comparison between the broad and narrow LVC sub-components.

TABLE 4.10: Mass loss rates for the LVC-BCs' using the method of Fang et al. (2018)

Target	V_c [km s ⁻¹]	V_{wind} [km s ⁻¹]	X_c [AU]	l_{wind} [AU]	$\text{Log}(L_{6300, LVC-BC})$ [L _⊙]	$\dot{M}_{wind}$ [M _⊙ yr ⁻¹]	η [%]
T = 5,000 K C(T) = 2.4×10^{-4}							
MP Mus	-0.04	-0.05	≤ 0.47	≤ 0.89	4.67	2.77×10^{-11}	0.55
GQ Lup	3.00	6.00	≤ 0.64	≤ 0.74	-4.41	7.63×10^{-9}	200.75
DG Tau	-53.00	-62.50	15.6	29.44	-3.74	9.20×10^{-9}	73.11
CW Tau	2.00	4.73	≤ 3.82	≤ 4.21	-4.06	2.37×10^{-9}	10.85
GK Tau	-4.00	-12.29	≤ 0.98	≤ 1.04	-5.33	1.32×10^{-9}	37.32
DK Tau	-5.00	-5.56	≤ 3.68	≤ 8.39	-4.71	3.11×10^{-10}	12.09
EX Lup	-15.00	19.04	≤ 3.95	≤ 6.42	-4.91	8.83×10^{-10}	33.56
T = 10,000 K C(T) = 2.6×10^{-5}							
MP Mus	-0.04	-0.05	≤ 0.47	≤ 0.89	4.67	3.00×10^{-12}	0.06
GQ Lup	3.00	6.00	≤ 0.64	≤ 0.74	-4.41	8.27×10^{-10}	21.75
DG Tau	-53.00	-62.50	15.6	29.44	-3.74	9.97×10^{-10}	7.92
CW Tau	2.00	4.73	≤ 3.82	≤ 4.21	-4.06	2.57×10^{-10}	1.18
GK Tau	-4.00	-12.29	≤ 0.98	≤ 1.04	-5.33	1.43×10^{-10}	4.04
DK Tau	-5.00	-5.56	≤ 3.68	≤ 8.39	-4.71	3.37×10^{-11}	1.31
EX Lup	-15.00	19.04	≤ 3.95	≤ 6.42	-4.91	9.56×10^{-11}	3.64

The absolute value of V_{wind} and l_{wind} are used to calculate the mass outflow rates. l_{wind} is the deprojection of the length measured from the 1D SA, X_c measured along the jet PA.

V_{wind} is the deprojection of the velocity, V_c measured from the kinematic fitting of the line profiles along the jet PA.

4.4 Discussion

Three questions that arise from the analysis presented above in the context of the current literature. Firstly, does the kinematic fitting reported here align with the fitting reported from the various studies examining large samples of CTTs? (e.g. Banzatti et al., 2019; Fang et al., 2018; Pascucci et al., 2020; Simon et al., 2016, etc.). If not, does this infer something about line profile variability in time and possibly the blending between different line sub-components? Two, what has the spectro-astrometric analysis above revealed about the nature of the LVC and does this align with what has been found for RU Lupi given the work of Whelan et al. (2021) and Birney et al. (2024b)? Three, do the outflow efficiency values derived above from the [O I] $\lambda 6300$ line align with the assumptions of current disc wind models or do they deviate as in the case of RU Lupi? Answering these questions will determine if the [O I] $\lambda 6300$ FEL is a "good" tracer of disc winds in CTTs.

TABLE 4.11: Mass loss rates for the LVC-NCs' using the method of Fang et al. (2018)

Target	V_c [km s ⁻¹]	V_{wind} [km s ⁻¹]	X_c [AU]	l_{wind} [AU]	$\text{Log}(L_{6300, LVC-NC})$ [L _⊙]	$\dot{M}_{wind}$ [M _⊙ yr ⁻¹]	η [%]
T = 5,000 K C(T) = 2.4×10^{-4}							
TW Hya	0.36	0.36	≤ 0.89	≤ 7.30	-5.32	5.74×10^{-12}	0.37
T Cha	1.00	2.79	≤ 0.61	≤ 0.65	-4.70/-4.75	$3.30 \times 10^{-9} / 2.10 \times 10^{-9}$	82.93/52.81
LkH α 330	-0.20	-0.27	≤ 2.64	≤ 3.95	-3.95	1.19×10^{-10}	1.69
GQ Lup	-14.00	-28.00	≤ 0.64	≤ 0.74	-5.28	4.80×10^{-9}	126.23
DG Tau	-10.00	-11.79	15.40	29.06	-4.66	2.13×10^{-10}	1.69
	-22.00	-25.94	22.60	42.65	-4.34	6.69×10^{-10}	5.32
CW Tau	4.00	9.46	≤ 3.82	≤ 4.21	-4.88	7.13×10^{-10}	3.26
GK Tau	-0.43	-1.32	≤ 0.98	≤ 1.04	-4.94	3.52×10^{-10}	9.91
DK Tau	-5.00	-5.56	≤ 3.68	≤ 8.39	-6.34	7.28×10^{-10}	0.28
EX Lup	-4.00	-5.08	≤ 3.95	≤ 6.42	-5.97	2.01×10^{-11}	0.77
T = 10,000 K C(T) = 2.6×10^{-5}							
TW Hya	0.36	0.36	≤ 0.89	≤ 7.30	-5.30	6.22×10^{-13}	0.04
T Cha	1.00	2.79	≤ 0.61	≤ 0.65	-4.70/-4.75	$3.58 \times 10^{-10} / 2.28 \times 10^{-10}$	8.98/5.72
LkH α 330	-0.20	-0.27	≤ 2.64	≤ 3.95	-3.95	1.29×10^{-11}	0.18
GQ Lup	-14.00	-28.00	≤ 0.64	≤ 0.74	-5.28	5.20×10^{-10}	13.68
DG Tau	-10.00	-11.79	15.40	29.06	-4.66	2.31×10^{-11}	0.18
	-22.00	-25.94	22.60	42.65	-4.34	7.25×10^{-11}	0.58
CW Tau	4.00	9.46	≤ 3.82	≤ 4.21	-4.88	7.73×10^{-11}	0.35
GK Tau	-0.43	-1.32	≤ 0.98	≤ 1.04	-4.94	3.81×10^{-11}	1.07
DK Tau	-5.00	-5.56	≤ 3.68	≤ 8.39	-6.34	7.89×10^{-13}	0.03
EX Lup	-4.00	-5.08	≤ 3.95	≤ 6.42	-5.97	2.18×10^{-12}	0.08

The absolute value of V_{wind} and l_{wind} are used to calculate the mass outflow rates. l_{wind} is the deprojection of the length measured from the 1D spectro-astrometry, X_c measured along the jet PA. V_{wind} is the deprojection of the velocity, V_c measured from the kinematic fitting of the line profiles along the jet PA.

4.4.1 Previously reported kinematics on this sample

Firstly, it is important to discuss and compare the kinematic fitting presented in Tables 4.3 and 4.4 to similar analysis available in the literature. There are at least five previous studies (Banzatti et al., 2019; Fang et al., 2018; Nisini et al., 2024; Pascucci et al., 2020; Simon et al., 2016) over the last decade that decompose the [O I] $\lambda 6300$ line profiles into its sub-components, two of which also cover the [S II] $\lambda 6731$ line also. However, before proceeding with this comparison it is important to note that each paper does not cover the same sample of targets so, the overlap with the sample of CTTs presented in this chapter varies between studies. Also, the velocity resolution of the spectra presented in these works and the spectra presented in this chapter vary. This impedes the comparison slightly as lower velocity resolution observations may be more subject to blending between different sub-components in each line profile. Given these caveats; the kinematic fitting is now discussed.

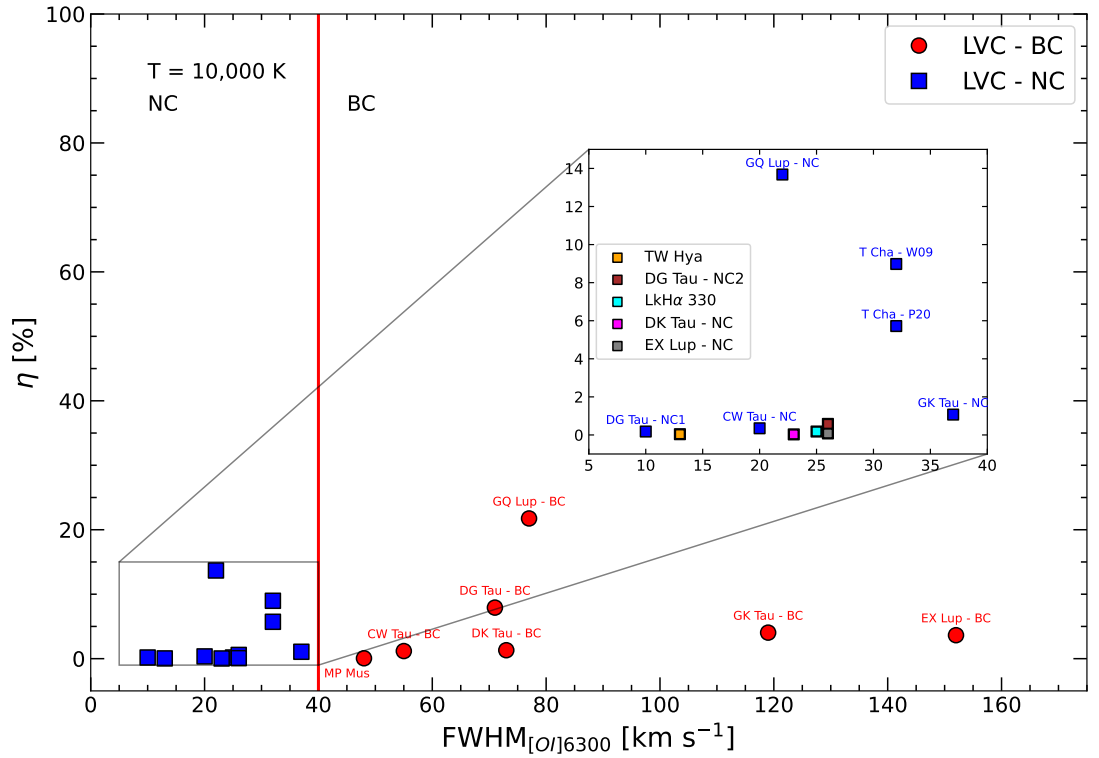


FIGURE 4.14: Efficiency of each [O I] $\lambda 6300$ line LVC sub-component at $T = 10,000$ K as a function of the FWHM of each LVC sub-component.

TW Hya

The [O I] $\lambda 6300$ line profile of TW Hya is straightforward, consisting of a single peak that is reproduced one narrow and slightly blueshift Gaussian component. The results presented here reflect this showing a peak velocity of 0 ± 0.3 km s⁻¹ and a FWHM of 13 ± 0.3 km s⁻¹. This line has also been detected by three other works (Banzatti et al., 2019; Fang et al., 2018; Simon et al., 2016) and has shown similar properties. Simon et al. (2016) report a peak velocity of -0.78 km s⁻¹ and a FWHM of 12 km s⁻¹. Fang et al. (2018) report a peak velocity of 0 km s⁻¹ with a FWHM of 12.8 km s⁻¹ and similarly Banzatti et al. (2019) purport that this profile has a peak velocity of -1.3 km s⁻¹ and a FWHM of 13 km s⁻¹. The result presented here generally align with the values reported in previous works. Also, a [S II] $\lambda 6731$ line profile is not reported in the previous literature and is not detected here as well.

T Cha

Pascucci et al. (2020) report the only other detection of the [O I] $\lambda 6300$ line profile in T Cha. They find a peak velocity of -0.4 km s^{-1} varying slightly from the value of $1 \pm 0.5 \text{ km s}^{-1}$ found from the fitting in this chapter. Additionally, the FWHM found by Pascucci et al. (2020) is 46.8 km s^{-1} which is considerably higher than the value of $32 \pm 0.5 \text{ km s}^{-1}$ found from the fitting presented in this chapter. Following the definition of LVC-NC and LVC-BC presented in Chapter 1, implies that this study detects the LVC of T Cha as narrow whereas, Pascucci et al. (2020), detects this component as broad. This could imply genuine variability in the [O I] $\lambda 6300$ line profile of T Cha between these two studies.

MP Mus

Like for T Cha, Pascucci et al., 2020 report the only other detection of a [O I] $\lambda 6300$ line in MP Mus. The value found here for the peak velocity is $0 \pm 0.4 \text{ km s}^{-1}$ which is approximately in agreement with the peak velocity of -0.5 km s^{-1} found by Pascucci et al. (2020). For the FWHM, the value found by Pascucci et al. (2020) (29.8 km s^{-1}) which is much smaller than the value found from the fitting presented in this study ($48 \pm 0.4 \text{ km s}^{-1}$). Given this, MP Mus may also be experiencing some variability in its [O I] $\lambda 6300$ line profile between these two works. MP Mus also does not have a detectable [S II] $\lambda 6731$ line.

LkH α 330

As was the case for the previous two stars, LkH α 330 has a single peaked [O I] $\lambda 6300$ line profile and no [S II] $\lambda 6731$ line. The only other detection of a [O I] $\lambda 6300$ line in this source is presented by Banzatti et al. (2019) where they fit a Gaussian of the same FWHM as what is presented in this study (i.e. 25 km s^{-1} in both works; with an uncertainty of $\pm 0.7 \text{ km s}^{-1}$ in this work) but, the peak velocity of this line profile is significantly more blueshifted in Banzatti et al. (2019) compared to this study (-5.4

km s^{-1} vs $0 \pm 0.7 \text{ km s}^{-1}$). As suggested before, that this change in the peak velocity between this study and Banzatti et al. (2019) may be a hint at variability in the [O I] $\lambda 6300$ line of LkH α 330 between these two observations.

GQ Lup

GQ Lup is the first MuC source in the sample with a line profile consisting of both an LVC-NC and LVC-BC. There has been sparse reporting on this source; with only two peak velocities of -2.9 km s^{-1} and -4 km s^{-1} reported by Pascucci et al. (2020) and Fang et al. (2018) respectively. One FWHM value of 91.4 km s^{-1} was also reported by Fang et al. (2018). The kinematic fit presented here is interesting given that it shows the [O I] $\lambda 6300$ line profile of this star consists of two LVC sub-components, one LVC-NC at $-14 \pm 0.3 \text{ km s}^{-1}$ and one slightly redshifted LVC-BC at a peak velocity of $3 \pm 0.3 \text{ km s}^{-1}$. This is in contrast to the SC [O I] $\lambda 6300$ line profiles reported by Fang et al. (2018) and Pascucci et al. (2020). It also appears here that the [O I] $\lambda 6300$ line profile of GQ Lup has experienced variability in shape over time. Particularly considering that the spectra presented in Fang et al. (2018) were taken on January and May 23 in 2008, over a decade before the spectra of GQ Lup were observed for this work.

DG Tau

DG Tau is a very interesting source. It is the only source in this sample of CTTs that has an [O I] $\lambda 6300$ line profile consisting of four sub-components. Additionally, it is the first star in this sample with a clear detection of [S II] $\lambda 6731$. These line profiles are observed to vary remarkably over time (see Otten et al., 2025, and Chapter 5 for more details). The [O I] $\lambda 6300$ line of DG Tau has been detected and characterised in a number of studies at various spectral resolutions (i.e. Simon et al., 2016 and Banzatti et al., 2019 at a resolution of $\sim 7 \text{ km s}^{-1}$, and Nisini et al., 2024 at a spectral resolution of $\sim 3 \text{ km s}^{-1}$). Out of these previous works, Nisini et al., 2024 presents the only other example of an [O I] $\lambda 6300$ line in the spectra of DG Tau with four sub-components. They infer that this is likely due to the higher spectral resolution of their spectra compared to others in the literature. However, the spectra presented above have a factor 2 lower

spectra resolution compared to Nisini et al. (2024) and still show four sub-components in the [O I] $\lambda 6300$ line of DG Tau. As such, it is pertinent to compare the profiles of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ lines from these two datasets. For ease of comparison, Table 4.12 compares the fit parameters of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ line fitting from this chapter and Nisini et al. (2024). Both this study and Nisini et al. (2024) measure the peak velocities and FWHM in the stellocentric rest frame

TABLE 4.12: Comparison of the kinematic fitting of the DG Tau [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs between this work and Nisini et al. (2024).

Line	Comp. Type	V_c [km s ⁻¹]	V_c - Nisini24 [km s ⁻¹]	FWHM [km s ⁻¹]	FWHM - Nisini24 [km s ⁻¹]
[O I] $\lambda 6300$	LVC-NC	-10	-3.9	10	12.3
	LVC-NC	-22	-16.9	26	24.0
	LVC-BC	-53	-60.0	71	78.0
	HVC	-141	-127.2	83	51.5
[S II] $\lambda 6731$	LVC-NC	-20	-17.1	27	28.3
	LVC-BC	-48	-50.4	60	69.6
	HVC	-134	-126.4	68	47.3

The [O I] $\lambda 6300$ of DG Tau shows two LVC-NCs in both studies. Both of the LVC-NCs and the jet component studied here appear to be more blueshifted than those observed in Nisini et al. (2024). The FWHM values of both LVC-NCs are comparable between both studies. The jet component observed here has a larger FWHM than in Nisini et al. (2024). The LVC-BC on the other hand is reported to be more blueshifted in Nisini et al. (2024) than the spectra presented here. The FWHM of the LVC-BC presented here is also smaller than the value reported by Nisini et al. (2024). For the [S II] $\lambda 6731$ line, the differences and similarities between this study and Nisini et al. (2024) appear the same over the line sub components. These variations may be attributed to blending between components in each of the line profiles due to the known slowing of the high velocity jet over the time period these two sets of observations were taken (Pyo et al., 2024; Takami et al., 2023). Additionally, there is a difference of ~ 4.7 km s⁻¹ between the measured heliocentric radial velocity in this study and Nisini et al. (2024) (19 km s⁻¹ here and 14.34 km s⁻¹ for Nisini et al. (2024)).

CW Tau

The [O I] $\lambda 6300$ line of CW Tau is detected in the studies of Banzatti et al. (2019) and Simon et al. (2016) and Nisini et al. (2024). In each of these studies both the LVC-NC and LVC-BC are slightly blueshifted whereas these components are slightly redshifted in the [O I] $\lambda 6300$ line presented here in Fig 4.5. The FWHM of the LVC-NC is comparable in all the published studies and in the fitting presented here. Although, the FWHM of the LVC-BC measured here is significantly lower than in the previously published works. It is noted that the jet component of the [O I] $\lambda 6300$ line profile shown here seems to have slowed (less blueshifted) compared to the previous studies. The slowing of the jet may imply that that jet may have partially blended in to the LVC-BC, reducing its FWHM.

The [S II] $\lambda 6731$ line of CW Tau was observed in the works of Simon et al. (2016) and Nisini et al. (2024). The results presented in this chapter and Nisini et al. (2024) agree that the observed LVC-NC in this line is blueshifted by less than 0.5 km s^{-1} . In opposition to this, Simon et al. (2016) reports a blueshift of $> 5 \text{ km s}^{-1}$. The FWHM values of this component are measured to be 24 km s^{-1} in both previous studies, whereas a FWHM of 33 km s^{-1} in the LVC-NC is measured in this study. The jet component of CW Tau in this line are observed to have comparable peak velocity in all three measurements (-111 km s^{-1} in this study, -116 km s^{-1} in Simon et al. (2016) and -116.7 km s^{-1} in Nisini et al. (2024)), although, it is noted here that the FWHM measured in this study for the jet component is larger than those measured by the other two studies.

GK Tau

Banzatti et al. (2019), Fang et al. (2018), and Simon et al. (2016) and Nisini et al. (2024) detect an [O I] $\lambda 6300$ line for this star. The main difference between the kinematic fitting presented here and the previous studies is the detection of a redshifted jet component in this line. This is interesting given that Banzatti et al. (2019) is the only other study to find a HVC, although they detect it as a blueshifted component. The [S II] $\lambda 6731$ line in this source has only previously been reported by Nisini et al.

(2024). The [S II] $\lambda 6731$ line profile in GK Tau varies dramatically between this work and Nisini et al. (2024). In this study, two HVCs are detected; one blueshifted and the other redshifted. Additionally, a LVC-BC is also observed. Nisini et al. (2024) finds only two components, an LVC-NC and a broad component at -69 km s^{-1} that they mark as a blue HVC.

DK Tau

The [O I] $\lambda 6300$ in DK Tau has been detected in four of the five quoted studies, excluding Pascucci et al. (2020). The fit presented here for the LVC-NC of this line agrees with the fits of Fang et al. (2018) and Banzatti et al. (2019) but disagrees with the fit of the LVC-NC presented by Simon et al. (2016). This may be attributed to possible line profile variability between observations. Nisini et al. (2024) do not report a LVC-NC in the [O I] $\lambda 6300$ line profile of DK Tau. The LVC-BC of this line varies between studies with the kinematic fit presented here in the closest agreement to the fit of Simon et al. (2016). The HVC fit in the [O I] $\lambda 6300$ profile of this star seems more blueshifted in the earlier studies of Fang et al. (2018) and Simon et al. (2016) and Banzatti et al. (2019) but, it appears to slow according to the fit of Nisini et al. (2024) and the kinematic fit presented in this chapter.

DK Tau has a [S II] $\lambda 6731$ line as reported here and in the studies of Nisini et al. (2024) and Simon et al. (2016). These previous studies report a very different line profile shape than what is reported here. Nisini et al. (2024) reports that the [S II] $\lambda 6731$ line of DK Tau consists of one HVC at a blueshifted velocity of greater than -130 km s^{-1} and Simon et al. (2016) purport a profile consisting of one narrow component at a peak velocity -56 km s^{-1} . These drastic differences seem to imply that variability of the [S II] $\lambda 6731$ line in DK Tau is significant.

EX Lup

Unlike DG Tau, CW Tau, GK Tau and DK Tau, EX Lup has only a marginal detection of [S II] $\lambda 6731$ line, where only a HVC is observed (see upcoming paper by Unni et

al. (in prep)). The shape and intensity of the [O I] $\lambda 6300$ line varies hugely between different studies. The shape and intensity of different line subcomponents seem to vary depending on if this star is in an out-bursting or quiescent state. This thesis reports that the LVC of EX Lup in [O I] $\lambda 6300$ consists of a slightly blueshifted LVC-NC and a more blueshifted LVC-BC.

4.4.2 The spectro-astrometric analysis in the context of RU Lupi

This subsection provides a discussion of the spectro-astrometric analysis shown above in the context of the pilot studies done on the CTTS RU Lupi (Birney et al., 2024b; Whelan et al., 2021). The combined findings of Whelan et al. (2021) and Birney et al. (2024b) indicate that the LVC-NC component of the [O I] $\lambda 6300$ line of RU Lup originates from within the gravitational potential well of the star and traces an MHD disc wind. This MHD disc wind was found to extract minimal excess angular momentum from RU Lupi star-disc system as the efficiency of the LVC-NC component in RU Lupi was found to be $\ll 1\%$ at either of the assumed gas temperatures of 5,000 or 10,000 K when applying the method of Fang et al. (2018).

The results presented in Tables 4.10 and 4.11 and in Fig. 4.14 imply that the mass loss rates and efficiencies of the LVC-NC and LVC-BC sub components change substantially between each CTTS and between each sub component. The efficiency of the LVC-NC ranges from 0.03% in the case of DK Tau, to over 13% for GQ Lup at the assumed gas temperature of 10,000K. For the LVC-BC, the efficiencies range from 0.06% for MP Mus, to 21.75% again for GQ Lup at $T = 10,000\text{K}$. At 5,000K, the mass loss rates increase by an order of magnitude and the efficiency values increase significantly compared to the values derived at $T = 10,000\text{K}$. The larger efficiency of the LVC-BC compared to the LVC-NC aligns with the interpretation of Banzatti et al. (2019), Pascucci et al. (2023), and Pascucci et al. (2020) that the LVC-BC is removing more excess angular momentum than the LVC-NC from the star-disc system. These studies suggest that this could be explained if the LVC-BC is less vertically extended above the disc

mid-plane than the LVC-NC.

The combination of the spectro-astrometric results shown in Figs. 4.7 to 4.13 and the mass flux/efficiency analysis shown by Fig. 4.14 seem to indicate two findings. One, the LVC-NC in RU Lup is likely unique in tracing an MHD disc wind. In most of the objects studied in this sample, the spectro-astrometric analysis of the LVC indicates that the LVC is spatially compact. The compact nature of the LVC is well known. Previous kinematic and previous studies have shown this (see Pascucci et al., 2023, and references therein). In addition, narrow field mode Multi Unit Spectroscopic Explorer (MUSE) observations of TW Hya have shown evidence that 80% of the [O I] $\lambda 6300$ emission from TW Hya is contained within 1 AU from the star (Fang et al., 2023b). Although, the [O I] $\lambda 6300$ has been observed to be spatially extended (HVC, tracing jets) for several other CTTs (Flores-Rivera et al., 2023). Two, the efficiency and mass loss rates of the LVC sub components varies significantly with temperature. The uncertainty on the temperature of the material traced by [O I] $\lambda 6300$ puts into question the accuracy of efficiencies and mass loss rates determined via the method of Fang et al. (2018).

Another caveat to this analysis in the [O I] $\lambda 6300$ line is the blending between different velocity components in the line profiles due to the limitations of the spectral resolution. This adds to the difficulty of separating the centroid offsets between the LVC sub-components. The spectro-astrometric analysis of the [S II] $\lambda 6731$ line in the LVC is minimal as the continuum subtraction routine essentially removes the ability to reliably fit the spatial profiles in this line region. The HVC spectro-astrometric analysis in both the [S II] $\lambda 6731$ and [O I] $\lambda 6300$ lines are also presented, but its interpretation is beyond the scope of this chapter and will be part of a future work. DG Tau will be discussed in more detail in Chapter 5.

4.4.3 Is [O I] $\lambda 6300$ a reliable wind tracer?

The [O I] $\lambda 6300$ is clearly useful in detecting low velocity emission in many of the CTTs studied to date. Therefore, it is a useful wind tracer. However, the results presented in this chapter suggest it is not a reliable line for testing wind efficiency or measuring the mass flux (and hence, angular momentum) extracted through a given wind. This section now outlines the reasoning for this.

1. As discussed in the last section, the uncertainty in the temperature of the gas traced by [O I] $\lambda 6300$ limits the accuracy to which the mass flux in the wind can be estimated. Nisini et al. (2024) reiterate this and imply that the LVC of [O I] $\lambda 6300$ may be significantly contaminated with emission from the upper disc atmosphere.
2. In addition, even with the sub-seeing spatial information extracted with SA, it is clear that the LVC in this sample of CTTs is compact in nature, regardless of source properties. This appears to imply that the emission is very close to the source. Hence, even with SA, the geometry of the emission traced by the [O I] $\lambda 6300$ line difficult to constrain. Given this, and the previous points made, Nisini et al. (2024) argue that the mass loss rates derived for the wind with the [O I] $\lambda 6300$ line using the method of Fang et al. (2018) may be off by a factor of up to two orders of magnitude.
3. In the single source (RU Lupi) where [O I] $\lambda 6300$ LVC is spatially resolved with spectro-imaging (Birney et al., 2024b). The derived mass flux in the [O I] $\lambda 6300$ LVC is extremely low and indicates that the [O I] $\lambda 6300$ line may be tracing a transitional layer in the outflow structure between the atomic part of the flow and the molecular flow. This "onion like" radial stratification structure in outflows has been observed on several occasions in both atomic and molecular emission lines over the last two decades (e.g. Agra-Amboage et al., 2014; Bacciotti et al., 2025; Lavalley-Fouquet, Cabrit, and Dougados, 2000; López-Vázquez et al., 2024; Pascucci et al., 2024). The results here indicate that the [O I] $\lambda 6300$ LVC does not constrain the mass flux well, as it varies significantly between stars. Blending

between line sub components may cause inaccuracies in the derived mass flux when using the [O I] $\lambda 6300$ line and uncertainty in the temperature exacerbates the uncertainty in the mass flux measurements. However, as stated previously, the [O I] $\lambda 6300$ itself is not always compact and may trace high velocity emission (Flores-Rivera et al., 2023). This work hence reiterates that the LVC is compact and the poor constraint on the mass flux limits the ability of the [O I] $\lambda 6300$ line to test for wind driven accretion.

4. The previous point regarding RU Lupi is further reinforced given the recent spectro-imaging observations of Pascucci et al. (2024). These observations have resolved a low velocity and wide wind structure in H_2 , surrounding a high velocity jet traced by [Fe II] in a sample of four CTTs (see their Fig. 2). Given this, the [O I] $\lambda 6300$ line seems to provide limited information about the mass flux of low velocity winds in these systems. Cooler wind tracers such as H_2 in the infrared could shed more light on this problem. Additionally, outflows traced by transitions of CO in the sub-mm (e.g. ^{12}CO) appear to trace low velocity winds with high mass flux. This could be explained if an MHD wind is extracting a large portion of the excess angular momentum contained within the star-disc system (de Valon et al., 2022).

4.5 Conclusions and Summary

In conclusion, the kinematics of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FEL profiles appears to vary significantly between this work and other studies over various periods of observations. This could be genuine variability of the line profiles over time however, there is a caveat, as there is very likely to be blending between different velocity components in the line profiles; particularly in [O I] $\lambda 6300$. This may be remedied by a multi epoch, high spectral resolution and high time cadence campaign to examine if these variations in the line profiles over time are real or due to some other effect.

The spectro-astrometric analysis of the LVC in [O I] $\lambda 6300$ shows that in most of the CTTSs presented in this sample (excluding DG Tau), the LVC appears to be very spatially compact. The LVC region of DG Tau will be further explored with multi epoch observations, discussed in Chapter 5. An offset in the LVC components of the [S II] $\lambda 6731$ line is seen for DG Tau and DK Tau where the centroid offset measured in this line is larger than for [O I] $\lambda 6300$. This is because the [S II] $\lambda 6731$ line has the lower critical density compared to [O I] $\lambda 6300$ (Chou et al., 2025; Giannini et al., 2019). It is noted that the centroid offset in the LVC/MVC of DK Tau looks blended with the HVC in both lines, so the offset measured may not be partly or wholly attributed to the LVC emission in both FELs. No offset is observed in LVC part of the [S II] $\lambda 6731$ line in CW Tau or GK Tau. The LVC region of this line in both of these stars appear very dim. Following the continuum subtraction routine, the spatial profiles of the LVC region of the [S II] $\lambda 6731$ line in both of these stars are incredibly dim. This makes spectro-astrometric analysis on the LVC emission region in this line for both stars challenging. Little information could be obtained after continuum subtraction of the [S II] $\lambda 6731$ line in these stars but, the non continuum subtracted position spectra of the LVC region of [S II] $\lambda 6731$ in both stars hints that this region is compact as in [O I] $\lambda 6300$.

The [O I] $\lambda 6300$ line appears to regularly trace low velocity emission. The work presented in this chapter, appears to indicate that the [O I] $\lambda 6300$ is limited in its ability to test the assumption that the wind mass loss rate is comparable to the stellar mass accretion rate. Several factors limit the [O I] $\lambda 6300$ in this endeavour. The first is the uncertainty on the temperature of the material being traced by this line. The second is blending between different LVC sub components. Thirdly, is the compactness of the LVC and difficulty SA has in spatially separating the LVC sub components in some of the CTTSs in this sample. The wind mass loss rates and efficiencies derived in this chapter using the method of (Fang et al., 2018) vary by an order of magnitude depending on the chosen temperature, and also vary between each CTTS in the sample. These uncertainties on the temperature of the material traced by the [O I] $\lambda 6300$ LVC, blending, and the possibility that the compactness of the LVC could imply that it is

contaminated with emission from the upper disc (Nisini et al., 2024). This limits the effectiveness of the [O I] λ 6300 line in tracing how low velocity winds extract angular momentum from star-disc systems

Additionally, the ability of the [O I] λ 6300 line to determine the efficiencies and mass fluxes from this low velocity emission does not appear to be dependent on disc inclination, (as the velocities and estimated wind heights were deprojected to account for this), mass or disc type (as three transition disc objects were included in this sample).

As such, it seems that the LVC of RU Lupi is unique in tracing an MHD disc wind. Cooler tracers such as H₂ lines and CO lines may provide more information regarding the mass flux and outflow efficiencies and in these slow winds. The additional information obtained via the study of colder wind tracers could put more context to the results obtained with the [O I] λ 6300 line from this study. Spectro-imaging and SA of H₂ and interferometric observations of CO in the sub-mm would provide more spatial information (e.g. collimation and the true wind height rather than the lower limit derived from spectro-astrometric analysis) and possible constraints on the temperature of the material within the low velocity outflow.

Chapter 5

Variability of the Low Velocity Component in the Forbidden Emission Lines of DG Tau

5.1 Chapter Introduction

The aim of this chapter is similar to that of Chapter 4 in that SA is used to understand the origin of the LVC in optical FELs. Where this chapter differs from the previous chapter is in the approach. Chapter 4 focused on the study of one epoch of observations in a small sample of Class II YSOs to examine the optical LVC. This chapter focuses on only one CTTS, DG Tau. Using multiple epochs observations, the aim is to study the variability in the LVC over time to see what can be ascertained about its origin. The star DG Tau was selected as its high velocity jet is known to have slowed over the last ~ 20 years (Takami et al., 2023). Many CTTS jets show kinematic variability which could be connected to variable accretion (Pyo et al., 2024). If the LVC traces an MHD wind then it is reasonable to expect that it will vary with a variable jet. The disc wind model for jet launching purports that the jet is the inner part of the larger MHD disc wind which becomes collimated by magnetic fields (Ferreira, Dougados, and Cabrit, 2006). Therefore, additional information on the origin of the LVC could be gleamed by studying variability. Variability of the LVC has been noted previously in the literature (Nisini et al., 2024; Simon et al., 2016) however, no dedicated study of the variability of the LVC has been conducted to date. This chapter provides the first dedicated study of variation in the LVC over time.

Three epochs of high spectral resolution data has been collected spanning the ~ 20 years over which the jet of DG Tau has slowed. These three epochs were observed in

2003, 2010 and 2021. Using these data, kinematic and spectro-astrometric analysis was performed to examine changes in the kinematic and spatial properties of the different sub-components of three physically distinct optical FELs ([O I] $\lambda 5577$, [O I] $\lambda 6300$ and [S II] $\lambda 6731$) over these epochs of time. These lines were selected because they have very different critical densities, decreasing from $\sim 10^8 \text{ cm}^{-3}$ for [O I] $\lambda 5577$ to $\sim 10^4 \text{ cm}^{-3}$ for [S II] $\lambda 6731$, hence, they trace different parts of the outflow system (Giannini et al., 2019).

The recent work of Chou et al. (2025) previously analysed the first epoch of high spectral resolution data of the FEL regions of DG Tau included in this study using kinematic fitting and spectro-astrometry. They identified seven distinct velocity components, six of which are associated with the DG Tau blueshifted outflow. They found two HVCs (HVC1, HVC2), a medium velocity component (MVC) and three low velocity components, (LVC-High, LVC-Medium and LVC-Low) for high, medium and low velocities within the LVC. Hereafter, these LVC sub-components are referred to as the LVC-H, LVC-M and LVC-L respectively. Table 5.1 provides explicit velocity ranges for the LVC-H, LVC-M and LVC-L as defined by the velocities at which the peaks and humps in the line profiles presented in Fig. 5.4 are identified.

TABLE 5.1: Velocity ranges for the LVC sub-components identified in this chapter.

Component	Velocity range ^a (km s^{-1})
LVC-H	$-58 < v < -31$
LVC-M	$-31 < v < -10$
LVC-L	$-10 < v < 0$

^a The velocities at which the peaks and humps in the line profiles shown in Fig. 5.4 are identified.

They argue that the HVC1 and HVC2 trace the extended jet and a stationary shock at the jet base respectively. Chou et al. (2025) also use line ratios and modelling to conclude that the LVC-H, LVC-M and LVC-L are associated with an interaction region between the jet and disc wind, a disc wind and the upper disc atmosphere respectively. Specifically, they argue that the absence of the LVC-H in the [O I] $\lambda 5577$ line

implies that the LVC-H must not be associated to material close to the disc surface as in the electron density and temperature in these close disc regions are enhanced given the [O I] $\lambda 5577$ line has a higher critical density than [O I] $\lambda 6300$ ($\sim 1 \times 10^8$ and 1×10^6 for [O I] $\lambda 5577$ and [O I] $\lambda 6300$ respectively). Additionally, they argue that the LVC-M must be associated to a wide angled disc wind as their spectro-astrometric analysis presents the negative velocity gradient predicted for such disc winds (Pontoppidan, Blake, and Smette, 2011; Whelan et al., 2021). Lastly, Chou et al. (2025) attribute the LVC-L to the disc atmosphere due to this sub-component having the highest [O I] $\lambda 5577$ to [O I] $\lambda 6300$ ratio, hence the highest electron density and temperature of the LVC sub-components identified. Given this, the authors argue a disc atmosphere emission origin for the LVC-L.

Their data is the first epoch of data used for the time variability study presented in this chapter. As such, for consistency the velocity components observed in this chapter are referred to using the nomenclature of Chou et al. (2025).

Also, it is noted here that all velocities quoted in this chapter are in the stellocentric reference frame, applying the velocity corrections outlined in Chapter 2. This chapter expands on the work of Chou et al. (2025) by looking at the changes in time of the velocity components of the blueshifted outflow, over a period of ≈ 18 years using an additional two epochs of high spectral resolution data. The rest of this chapter is organised as follows. Section 5.2 outlines the observations undertaken and the data reduction and analysis steps. Section 5.3 presents the PV diagrams and the results of the kinematic and spectro-astrometric analysis. Section 5.4 provides a discussion on the findings of this analysis. Lastly, in section 5.5 a summary the findings of this chapter is given alongside conclusions.

5.2 Observations, data reduction & analysis

DG Tau was observed in 2003 with the HDS on SUBARU telescope (Chou et al., 2025) and in 2010/2021 with the UVES (Dekker et al., 2000) instrument on the VLT (see Table

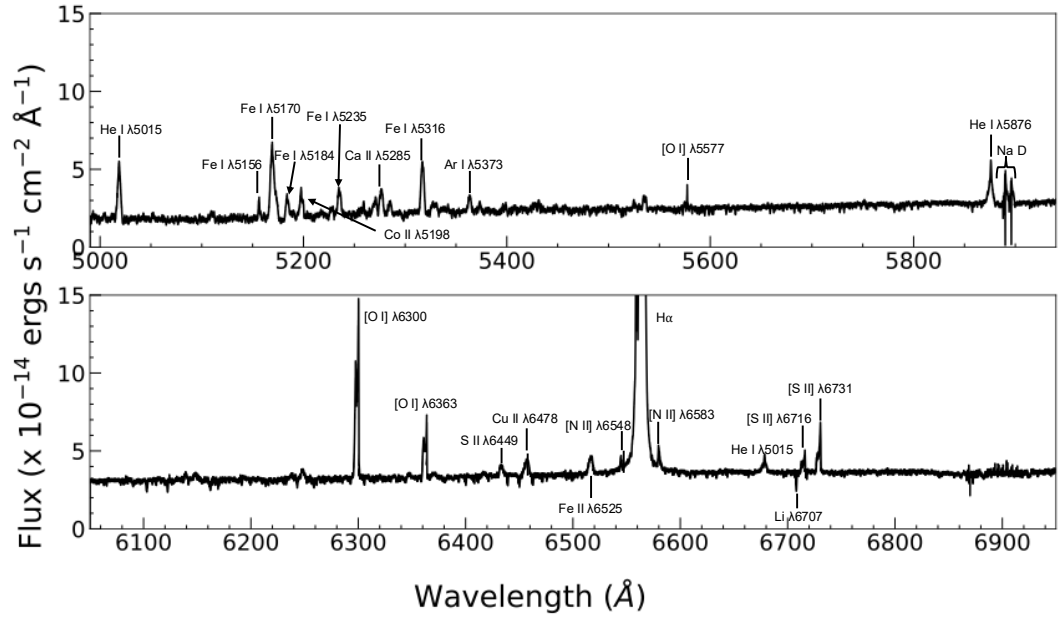


FIGURE 5.1: Full spectrum of DG Tau from 5000 to 7000 Angstroms along the axis of the micro-jet ($PA = 46^\circ$) in the topocentric rest frame. The most prominent emission and absorption features have been marked.

5.2). The details regarding the reduction of the 2003 data is presented in Chou et al. (2025). The UVES data were initially reduced using version 6.1.6 of the UVES pipeline (Freudling et al., 2013b). An example of a 1D flux calibrated UVES spectrum along the jet PA of DG Tau is presented in Fig. 5.1 and the physical properties of DG Tau relevant to this study are presented in Table 5.3. PV diagrams were constructed from the full spatial range of the 2D spectra generated from the UVES pipeline after continuum subtraction (see Chapter 2, Section 2.2.3 for details regarding continuum subtraction). These diagrams are shown for each line and epoch in Fig. 5.3.

The PV diagrams were constructed following a simple method. First, a background region of the spectrum was selected by eye (Fig. 5.2, red rectangle). This region only contained noise and no emission. From this region the rms noise and standard deviation was calculated for the background of each 2D spectrum.

A custom *Python* function was used to define the contour levels based on the values of

TABLE 5.2: Observation Log

Date (dd-mm-yyyy)	Slit PA (°)	Exposure Times (ncycles × s)
06-12-2003	44.4	4 × 900
06-12-2003	224.4	4 × 900
07-12-2003	134.4	4 × 600
07-12-2003	314.4	4 × 600
07-01-2010	46	2 × 755 ^a
14-11-2021	46	1 × 300
14-11-2021	136	1 × 300
16-11-2021	46	3 × 800
13-02-2022	136	3 × 800

^a Only one exposure was used as the source had moved in the slit during the second exposure.

TABLE 5.3: Properties of DG Tau.

Property	Value	Reference
SpTy	K7	Banzatti et al. (2019)
Mass ($M_{\odot}$)	0.80	Banzatti et al. (2019)
Radius ($R_{\odot}$)	1.85	Grankin (2016)
Age (yrs.)	3×10^5	Beckwith et al. (1990)
$\log(\dot{M}_{\text{acc}})$ ($M_{\odot}/\text{yr}$)	-8.39	Herczeg and Hillenbrand (2008)
Disc Type	full	Pascucci et al. (2015)
Disc PA (°)	119 ± 24	Isella, Carpenter, and Sargent (2010)
Blue Jet PA (°)	226 ± 1	Liu et al. (2016)
Disc incl. (°)	32 ± 2	Banzatti et al. (2019)
A_v (mag)	1.6 ± 0.15	Herczeg and Hillenbrand (2014)
Distance (pc)	125.3	GDR3 ^a
V_{RV} (km s^{-1})	19.0 ± 0.5	this work

^a GDR3 = Gaia Data Release 3 (Prusti et al., 2016; Vallenari et al., 2023).

the background rms noise and standard deviation (Fig. 5.2). For example, the first contour is defined as rms noise plus X times the standard deviation. The next level is just a scaling factor (e.g. $\sqrt{3}$) times the value of the first contour. Any subsequent contours are just defined by the scaling factor times the value of the previous contour. Using the FITS file header information, an assumed distance to DG Tau (Table 5.3 and the known rest wavelength of each FEL the axes of the of each PV diagram were converted into a spatial axis in AU and a velocity axis in the stellocentric reference frame. 1D spectral line profiles were then extracted from these 2D continuum subtracted spectra. The

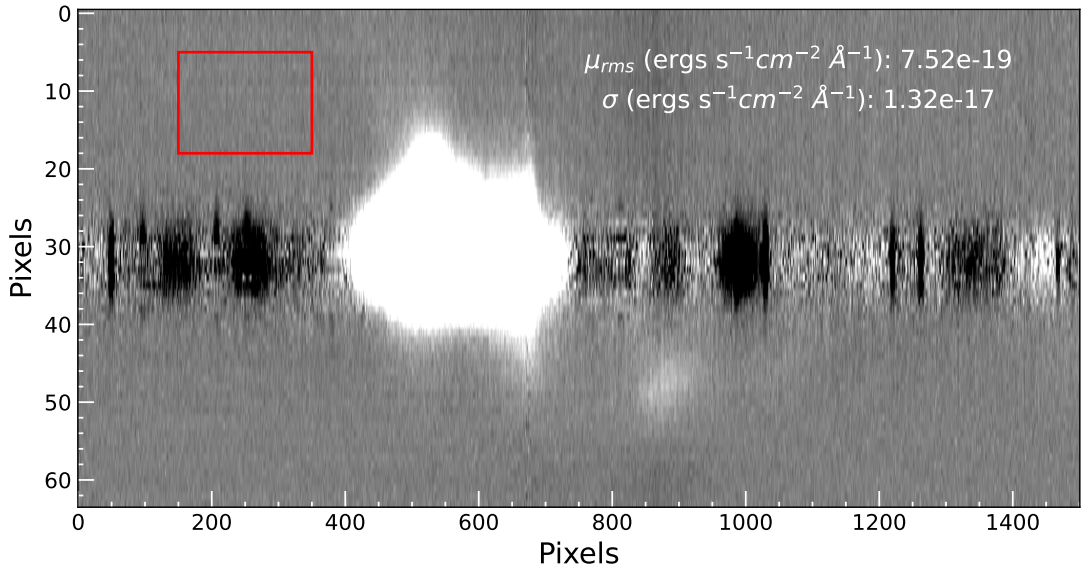


FIGURE 5.2: Example showing the method of how each PV diagram is constructed. A background region is selected (red rectangle) is selected by eye, from which the rms noise and standard deviation is calculated. The rms noise and standard deviation of the background is used to define the contour levels in each PV diagram. The example here shows the [O I] $\lambda 6300$ line region from DG Tau in 2021.

region in which the line profiles are extracted covers the whole spatial range of the 2D data. As shown in Chapter 2, the line profiles were then corrected for telluric and photospheric absorption features using the ESO MOLECFIT pipeline (Kausch et al., 2015; Smette et al., 2015) and the photospheric correction method first proposed by Hartigan et al. (1989). The kinematic fitting of the 1D spectra was then carried out following the procedure outlined in Chapter 4. In brief, this method involves generating a composite model of independent Gaussian components used to reproduce the shape of each line profile. The kinematic fitting to all FELs in each epoch is presented in Fig. 5.4 and Table 5.4. The spectro-astrometric analysis was performed on the [O I] $\lambda 5577$, [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs in each epoch of data for DG Tau in the fashion described in Section 2.1.

5.3 Results

Using PV diagrams and kinematic fitting procedure discussed in Chapter 4 (Table 5.4, Figs. 5.3 and 5.4) all the kinematic components first identified by Chou et al. (2025) are

TABLE 5.4: Kinematic fitting of the blueshifted outflow components in km s^{-1} . The centroid velocity and FWHM (in brackets) of each Gaussian are given for the HVC1, HVC2, MVC, LVC-H, LVC-M and LVC-L. The uncertainty on the centroid velocities and FWHM values is taken as 0.5 km s^{-1} . The spectra here shown in Fig. 5.4 and are taken from the full spatial range of the PV plots shown in Fig. 5.3. The extended jet component (i.e. the HVC1, marked in **bold**) for the [O I] $\lambda 6300$ line in 2003/2010 and the [S II] $\lambda 6731$ line in 2003 are identified as part of the blueshifted wing and not identified as a separate component by the kinematic fitting. To measure these values a further spectrum was extracted (see Fig. 5.5) from the extended jet region in 2003 and 2010.

	HVC 1	HVC 2	MVC	LVC-H	LVC-M	LVC-L
[O I] $\lambda 5577$						
2003	-	-191 (165)	-	-	-17 (23)	-
2010	-	-	-	-	-	-10 (21)
2021	-	-	-58 (198)	-	-	-12 (17)
[O I] $\lambda 6300$						
2003	-250 (50)	-185 (112)	-99 (67)	-46 (54)	-20 (21)	-9 (6)
2010	-178 (82)	-122 (126)	-62 (75)	-	-27 (41)	-8 (13)
2021	-141 (83)	-	-	-53 (71)	-22 (26)	-10 (10)
[S II] $\lambda 6731$						
2003	-249 (46)	-	-131 (207)	-53 (49)	-30 (19)	-
2010	-190 (26)	-	-97 (165)	-	-40 (49)	-
2021	-134 (68)	-	-	-48 (60)	-20 (27)	-

found but not in all lines or all years. Once identified, the blueshifted components of the lines in each epoch are investigated further using SA. The redshifted outflow was not studied due to its relative faintness. Only the blueshifted outflow was studied.

The differences between the emission line regions can be explained in terms of their different critical densities and the decreasing electron density of the outflow with distance from the star. [O I] $\lambda 5577$ traces the densest and therefore generally the most compact components. [O I] $\lambda 6300$ is the most complex with contributions from all the identified kinematic components. [S II] $\lambda 6731$ traces the most extended emission.

In Figs. 5.6 and 5.7 for example, the [S II] $\lambda 6731$ emission is extended up to twice the distance of the [O I] $\lambda 6300$ emission. The HVC2 and LVC-L are not identified in the [S II] $\lambda 6731$ line and Chou et al. (2025) argue that this is because the critical density of [S II] $\lambda 6731$ is too low. The effectiveness of the kinematic fitting for identifying the different velocity components is limited by the relative brightness of the components,

and it should be noted that a non-detection by the kinematic fitting does not mean that a particular component is not present in the outflow. For example, even though the LVC-L is not detected in the kinematic fitting of the [S II] $\lambda 6731$ line it does not mean that it is not present. Its contribution may be quenched due to critical density effects, but it may still contribute to the overall line profile as described further in Section 5.3.2. The results for the kinematic components associated with the jet by Chou et al. (2025) (HVC1, HVC2, MVC and LVC-H) and the LVC-L and LVC-M sub-components are now presented separately.

5.3.1 Variability in the jet components

As the HVC1 is part of the extended jet it can be studied through the PV diagrams and the kinematic fitting. It is always present in [O I] $\lambda 6300$ and [S II] $\lambda 6731$ and its decrease in velocity by $\sim 100 \text{ km s}^{-1}$ from 2003 to 2021 follows the known decrease in the jet velocity over this period (Pyo et al., 2024; Takami et al., 2023). In the 2003 and 2010, in the extracted spectra shown in Fig. 5.4, it appears as a blue-wing and is not picked up by the kinematic fitting. To measure the 2003 and 2010 HVC1 velocity and FWHM for Table 5.4, a second spectrum extracted from the jet region was used. This is presented in Fig. 5.5. The extended jet does not have a high enough excitation to produce detectable emission in [O I] $\lambda 5577$ with the exposure times and equipment used, hence this component is not detected in the kinematic fitting or PV plots of this emission line.

The HVC2 is detected in [O I] $\lambda 6300$ in 2003 and 2010, in both the PV plots and the kinematic fitting, but not in 2021. In 2021, it is likely blended with the HVC1 and MVC. The [S II] $\lambda 6731$ line does not trace this component and [O I] $\lambda 5577$ traces the HVC2 in 2003 only. The decrease in the HVC2 velocity from 2003 to 2010, is in line with the decrease in the HVC1 velocity.

The MVC and LVC-H are not always detected and only together in [O I] $\lambda 6300$ and [S II] $\lambda 6731$ in 2003. This is likely due to blending between the components as the velocity of the jet decreases. Chou et al. (2025) associate the LVC-H with a jet/wind

interaction region. They do not discuss the origin of the MVC but the decrease in velocity of this component from 2003 to 2021 may suggest that it is tracing newly ejected lower velocity knots which eventually become part of the extended jet.

Figure 5.6 presents the position spectra of the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ line regions in the three years. The [O I] $\lambda 5577$ line is not shown in this figure as the extended jet is faint or absent in all epochs of data for this line. The strong variability in the centroids beyond $\sim -75 \text{ km s}^{-1}$ trace the changes in the jet as the new knots are ejected with decreasing velocities. The region of the HVC1 initially shows a positive velocity gradient as material is accelerated in the jet (Whelan et al., 2021). The peaks and troughs in the region of the position spectrum that covers the jet emission are typical of knotty jets with the knots representing shocks in the outflow. For example, Fig. 5 of Whelan, Ray, and Davis (2004) and many of the figures presented in Hirth, Mundt, and Solf (1997) show similar signatures in their position spectra consistent with knotty jets. The HVC2 is located at $\sim 30 \text{ AU}$ in [O I] $\lambda 6300$ in 2010 in agreement with the Chou et al. (2025) analysis of the 2003 HDS spectra. The region of the MVC and LVC-H in both lines also has a positive velocity gradient, linking them to the jet, although in some cases the gradient is not steep.

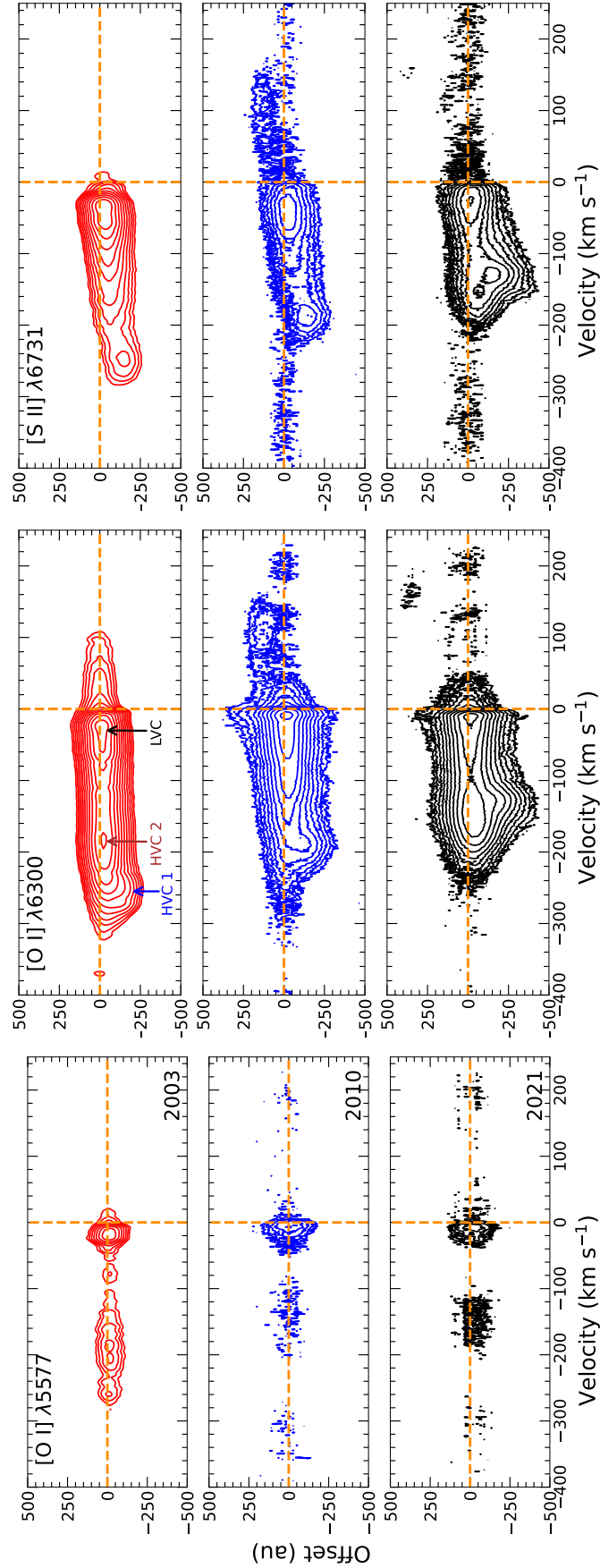


FIGURE 5.3: Position velocity diagrams of the [O I] $\lambda 5577$ (left) [O I] $\lambda 6300$ (middle) and [S II] $\lambda 6731$ (right) FELs in 2003, 2010 and 2021 from DG Tau. Contours start at 3σ (2003) and 7σ (2010, 2021) above the background rms noise and increase by factors of $\sqrt{2}$ (2003) and $\sqrt{3}$ (2010, 2021). The kinematic components identified by Chou et al. (2025) in the 2003 data (HVC1, HVC2 and LVC) are marked. The assumed distance to DG Tau is 125 pc (see Table 5.3).

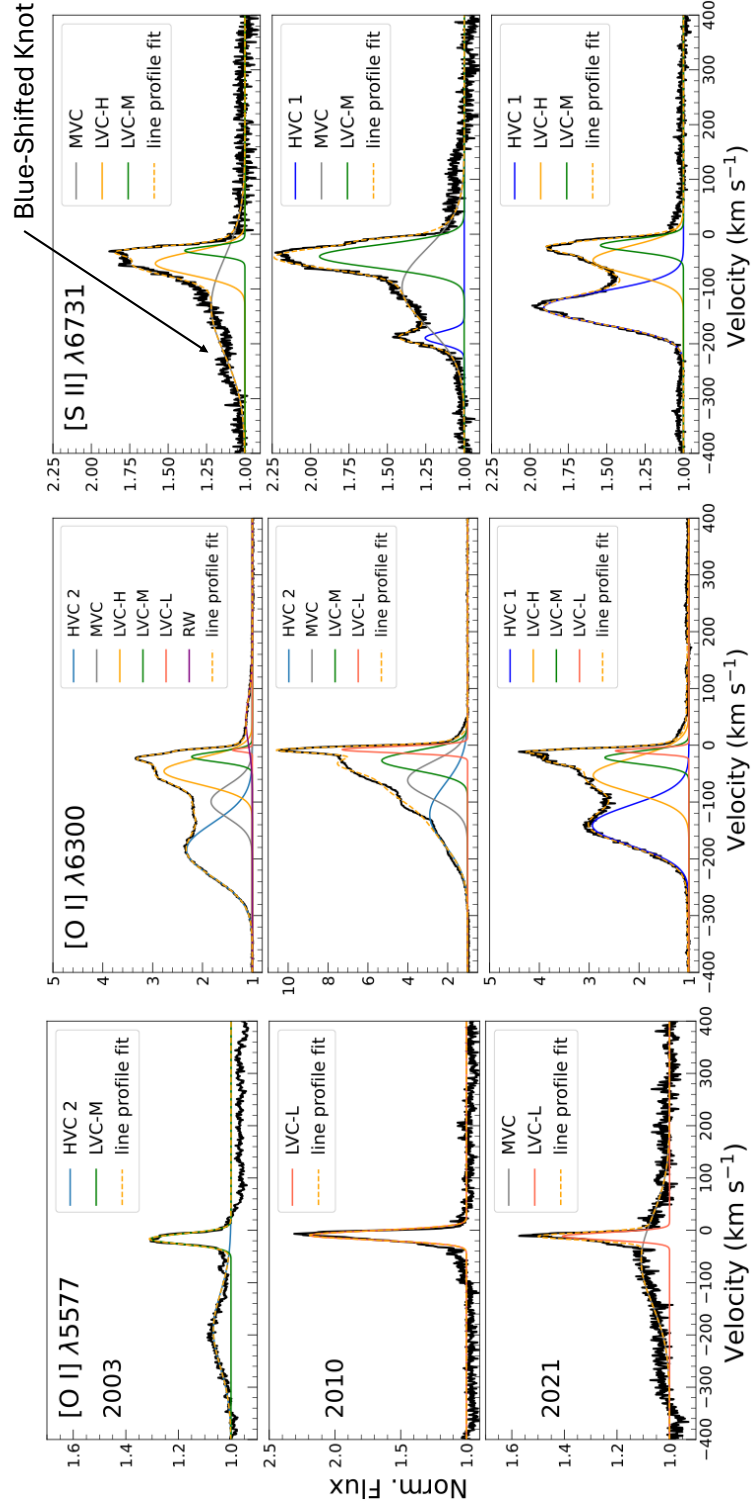


FIGURE 5.4: Kinematic fitting comparison of the [O I] $\lambda 5577$, [O I] $\lambda 6300$ and [S II] $\lambda 6731$ line profiles via Gaussian decomposition in 2003, 2010 and 2021. The line profiles are shown in the stellocentric frame and have been normalised to the continuum level. These profiles are extracted over the full spatial range of the PV diagrams shown in Fig. 5.3.

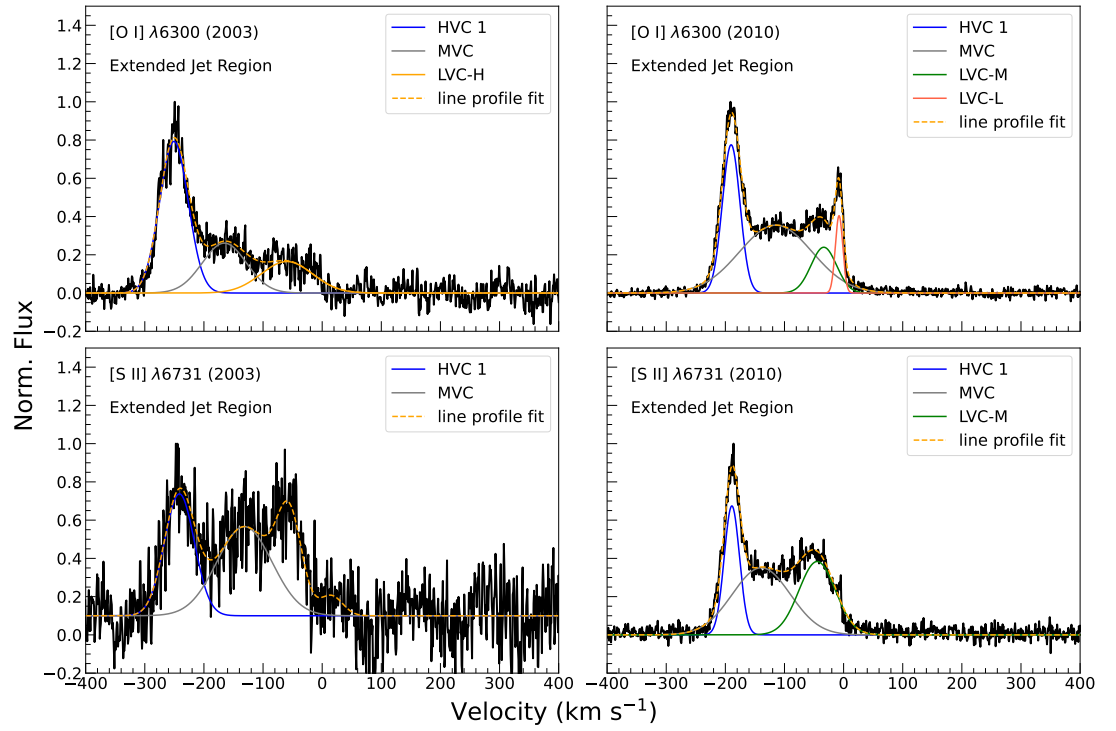


FIGURE 5.5: Kinematic fitting of the spectra extracted from the extended jet region of DG Tau in [O I] $\lambda 6300$ and [S II] $\lambda 6731$ in 2003 and 2010.

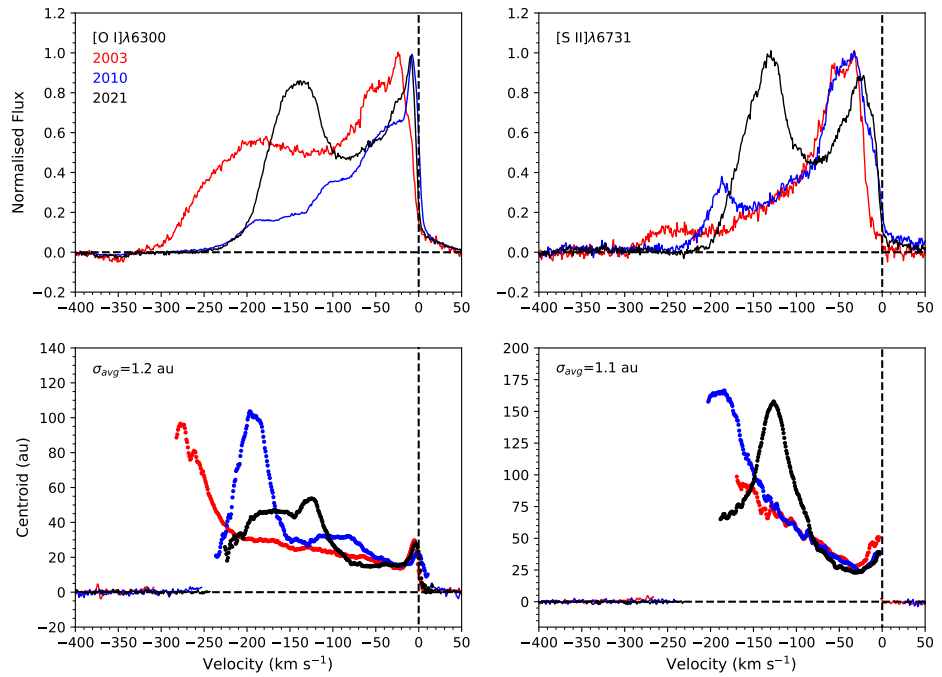


FIGURE 5.6: Top row: Comparison of the line profiles of [O I] $\lambda 6300$ and [S II] $\lambda 6731$ for each epoch of data. The line profiles have been normalised to max peak height to emphasise changes in the line shape over time. Bottom row: Position spectra for the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ lines in 2003 (red), 2010 (blue) and 2021 (black). The analysis of the [O I] $\lambda 5577$ line is not presented here as it does not show an extended jet component due to its high critical density.

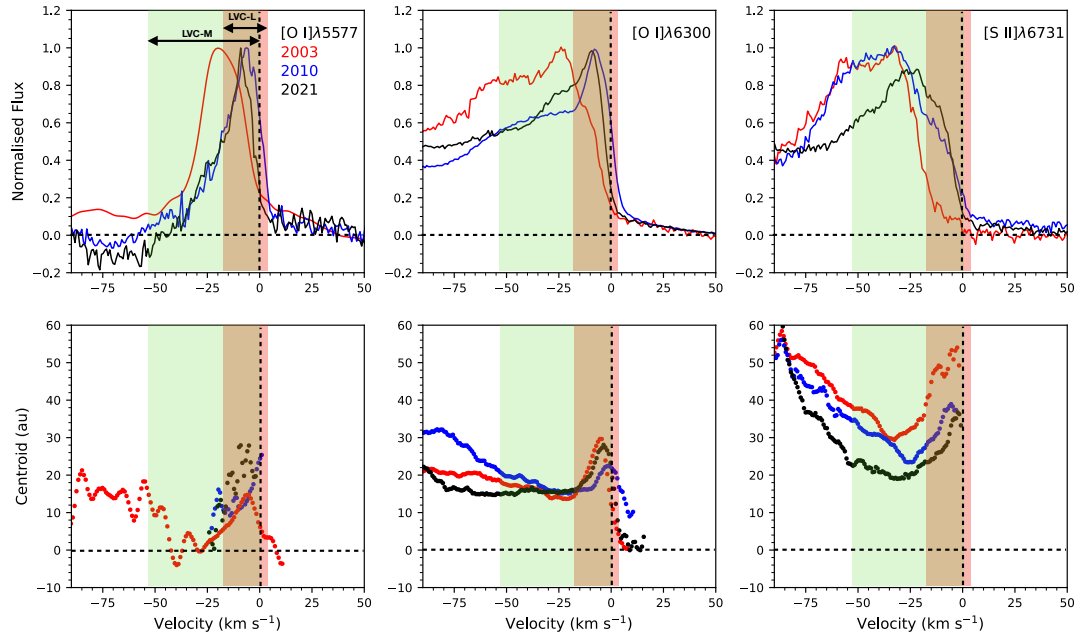


FIGURE 5.7: Top row: Comparison of the line profiles in the LVC region of each FEL in each epoch of data. 2003 = red, 2010 = blue and 2021 = black. These panels show a zoom of the line profiles shown in Fig. 5.6 from -90 to 50 km s^{-1} . This is done to emphasise the change of line shape in the LVC region over time. The line profiles have been normalised to max peak height. Bottom row: The position spectra of the LVC region for each FEL in each epoch of observation. 2003 = red, 2010 = blue and 2021 = black. The $1\text{-}\sigma$ uncertainties range from 1 to 5 AU where 5 AU is measured for the lowest S/N emission i.e. [O I] λ 5577 in 2021.

5.3.2 Variability in the LVC

The LVC-M and LVC-L kinematic components are now discussed in terms of changes in the shape of the line profile, the peak velocities measured by the kinematic fitting and the spatial properties of the emission regions as mapped with SA. These components are not spatially resolved in the PV diagrams of Fig. 5.3. The top row of Fig. 5.7 shows the LVC regions of each line profile with each epoch over-plotted. The lines have been normalised to the line peaks so changes in the LVC line profile shape become clear. As can be seen from Table 5.4, the peak velocities of these components do not vary greatly, unlike the jet components. The range of the LVC-M is marked in green in Fig. 5.7 and the LVC-L is marked in red, with the overlap region appearing in light brown. This brown colour highlights the strong overlap between the two components which impacts the kinematic fitting and SA. The line profile shape for all three lines changes substantially across the three epochs. This changes what components are detected by the kinematic fitting, the measured peak velocities and peak centroids measured by SA.

For [O I] $\lambda 5577$ in 2003 the line is at its widest and spans both the LVC-M and LVC-L. The LVC-M is very prominent in this epoch and therefore the kinematic fitting only detects a LVC-M. In 2010 and 2021 an LVC-L is detected by the kinematic fitting and the LVC-M is not, as what was the LVC-M has reduced to a blue-wing. The peak velocities recorded for these components are consistent with the other emission lines.

The [O I] $\lambda 6300$ LVC consists of a peak, an extended blue-wing and a less substantial red-wing. In 2003 the kinematic fitting measures the line peak at -20 km s^{-1} and this is identified as the LVC-M and a hump in the red-shifted wing is measured at -9 km s^{-1} and identified as the LVC-L. The shape of the line profile changes substantially in 2010 and 2021. The LVC-M is associated with a blue-wing in 2010 and a hump in the blue-wing in 2021, rather than a clear peak as in 2003. The LVC-L which appears as a hump in the red wing in 2003 is a clear line peak in 2010 and 2021. The peak velocity of the LVC-M increases from -20 km s^{-1} in 2003 to -27 km s^{-1} in 2010 and back to

-22 km s^{-1} in 2021. From this, it is argued that this velocity variation for the LVC-M in $[\text{O I}] \lambda 6300$ is due to some blending with the MVC jet component which is prominent in 2010 and the other changes to the line profiles described above. The peak velocity of the LVC-L does not vary by more than 2 km s^{-1} .

The shape of the $[\text{S II}] \lambda 6731$ LVC also changes between 2003 and 2021. The red-wing seen in 2003 becomes a hump in 2010 and 2021. It may be that the LVC-L is contributing to the line profile in 2010/2021 but, this contribution is not picked up by the kinematic fitting. This interpretation to explain the changes observed in the $[\text{S II}] \lambda 6731$ LVC aligns with the increase in the contribution of the LVC-L to both $[\text{O I}]$ lines in those years. The strong and wide blueshifted emission peak seen in 2003 and 2010 reduces to a blue-wing in 2021. The peak velocity of the LVC-M increases from -30 km s^{-1} in 2003 to -40 km s^{-1} in 2010 and back to -20 km s^{-1} in 2021. Again, it is argued that these changes to the peak velocity of the LVC-M is due to blending with the MVC jet component which is prominent in 2010 and the other changes to the line profiles described above.

The bottom row of Fig. 5.7 presents the position spectra of the LVC-M and LVC-L regions, for all three FELs, and in all three epochs, after continuum subtraction. Strikingly while as highlighted above, the shape of the LVC profiles change greatly, the position spectra in this region are very consistent across the nine measurements. Particularly when considering the velocities at which the peak centroids are measured and the velocity gradients. The biggest change in the velocity at which the maximum centroid is measured ($\sim -4 \text{ km s}^{-1}$) occurs in the $[\text{O I}] \lambda 6300$ emission region where both the LVC-M and LVC-L have comparable contributions. In contrast, no such change in the velocity of the maximum offset is observed for the $[\text{S II}] \lambda 6731$ measurements, in which the contribution of the LVC-L is minimised. For $[\text{O I}] \lambda 5577$, the position spectrum is most like what is seen in $[\text{O I}] \lambda 6300$ and $[\text{S II}] \lambda 6731$ in 2003, when the $[\text{O I}] \lambda 5577$ emission is the least noisy.

The velocity gradients are also remarkably consistent. A positive gradient is seen for all lines at velocities more blueshifted than -25 km s^{-1} and it marks where we begin to see blending between the LVC and jet components. A negative velocity gradient where offset increases with decreasing velocity is indicative of a disc wind and is observed in all emission lines and in all epochs for the region covering 0 to -25 km s^{-1} .

Where variability in the position spectra is observed is in the maximum and minimum values of the spatial centroids. If the LVC-L is tracing the upper disc atmosphere it would be more compact than the extended disc wind traced by the LVC-M. As the components are blended, the measured centroids are a combination of the position of both components. Where contributions from the LVC-M and LVC-L to the overall line profile are comparable, as in $[\text{O I}] \lambda 6300$, there is not much change in the maximum centroid between the years. In $[\text{S II}] \lambda 6731$ the peak centroid decreases from $\sim 55 \text{ AU}$ in 2003 to $\sim 35 \text{ AU}$ in 2010/2021. This is further evidence that the change in the blue-wing of the $[\text{S II}] \lambda 6731$ emission in those years is due to an increased contribution from the LVC-L. The peak centroid is less in 2010/2021 as it includes a stronger contribution from the more compact LVC-L.

Also, it is noted that the spatial centroids of the $[\text{O I}] \lambda 5577$ LVC-M reach a minimum value of 2 AU at -25 km s^{-1} in 2003 while in $[\text{O I}] \lambda 6300$ they are $\sim 15 \text{ AU}$ in all years and vary between 20 AU and 32 AU in $[\text{S II}] \lambda 6731$. This is evidence that the jet components are combining with the LVC-M to increase the measured offset of the more compact LVC-M emission in a similar way to how the LVC-L decreases it at velocities close to 0 km s^{-1} . There is no jet emission in $[\text{O I}] \lambda 5577$ in 2003 and so this does not occur in this case. The position spectra of the $[\text{O I}] \lambda 5577$ line in 2010 and 2021 are more difficult to interpret due to the relative faintness of the emission in those years so no conclusions are made. Similarly, no conclusions are made about the position spectra of the LVC red wing as measurements are impacted here by the accuracy of the continuum subtraction.

5.4 Discussion

The results for the jet components reveal remarkable spectral and spatial variability which is connected to the slowing of the DG Tau jet. The HVC1 is tracing a jet. This component decreases in velocity from 2003 to 2021. Pyo et al. (2024) correlate the changes in the velocity of the jet with corresponding changes in stellar mass accretion rate. The variability of the stellar mass accretion rate alongside hot spots on the stellar surface was also found to be linked to variability in the photometric magnitude of DG Tau. Pyo et al. (2024) argue that the variable stellar accretion causes an increase in the radius at which the high velocity jet is launched and thus a decrease in the jet velocity is seen. This is because the jet velocity is approximately proportional to the Keplerian velocity at the launch radius of the jet (Pudritz et al., 2007; Shu et al., 2000).

Chou et al. (2025) conclude that both the HVC1 and HVC2 components of the DG Tau FELs originate from the postshock regions in the jet, supported by the fact that these components are observed in emission lines from significantly different ionization conditions e.g. [O I] and He I. They measure the density of the HVC2 at $n_e \sim 10^6 \text{ cm}^{-3}$, which is higher than the HVC1 ($n_e \sim 10^4 \text{ cm}^{-3}$) and find using SA that it is located at $\sim 0''.2$ ($\sim 25 \text{ AU}$) from DG Tau. This is all strong evidence that it is the base of the DG Tau jet and is associated with a stationary shock component previously detected in X-rays (Güdel et al., 2011). The HVC2 is found in this study in the 2010 UVES spectra at a similar distance but at a lower velocity, $\sim -122 \text{ km s}^{-1}$ compared to -185 km s^{-1} in 2003. This is further evidence that it is a shock at the base of the jet, as the jet also decreases in radial velocity by $\sim 70 \text{ km s}^{-1}$ over this time period. Güdel et al. (2011) also report the emission from the X-ray knot to be variable which could explain its non-detection in [O I] in some years in this work.

The MVC is another velocity component associated with the jet. Its connection to the jet is supported by the decrease in its radial velocity from 2003 to 2021 in both [O I] $\lambda 6300$ and [S II] $\lambda 6731$ and the positive velocity gradient recorded by the spectro-astrometric analysis. This study proposes that the MVC is emission from the new

lower velocity knots being ejected into the DG Tau blueshifted jet and this could be confirmed by additional epochs of data. The non-detection of this component in [O I] $\lambda 5577$ is extra evidence that it is a shock in the jet.

The LVC-H is the final jet component to consider. In this study and that of Chou et al. (2025) the LVC-H is not detected in [O I] $\lambda 5577$ which points to it having a lower density and/or temperature than the components detected in [O I] $\lambda 6300$ and thus it is likely not coming from a region close to the star. Chou et al. (2025) discuss how gas from the wind entrained by the fast jet could explain the LVC-H and reference the detection of a similar component in the [Fe II] $1.64 \mu\text{m}$ emission from DG Tau studied by Pyo et al. (2003). In the observations of Pyo et al. (2003) the offset of the entrained component increases with velocity from $\sim -60 \text{ km s}^{-1}$ and at $\sim 30 \text{ AU}$ from the star. This is in agreement with the velocities of the LVC-H, the spatial centroids and the positive velocity gradient recorded in this study.

While the variability in the jet components is significant due to new ejections in the jet at decreasing velocities, the changes observed in the LVC are primarily driven by varying contributions from the LVC-M and LVC-L components, which mainly affect the shape of the line profile. Figure 5.7 shows that the contribution from the LVC-L to the overall line shape increases from 2003 to 2021 in all lines, while the contribution from the LVC-M decreases over that period. The contribution from the LVC-L to the [S II] $\lambda 6731$ line is never strong enough for it to be detected through the kinematic fitting.

Chou et al. (2025) argue that the LVC-L comes from the upper disc atmosphere based on the LVC-L [O I] $\lambda 5577$ to [O I] $\lambda 6300$ line ratio. This component has the highest value for this ratio indicating that it has the highest density or temperature (see their Fig. 7). Their evidence for the origin of the LVC-M in a disc wind comes from the negative velocity gradient recorded between 0 to -17 km s^{-1} for their analysis of the [O I] $\lambda 5577$, [O I] $\lambda 6300$ and [S II] $\lambda 6731$ lines. This is a strong signature of a disc wind.

As outlined in Whelan et al. (2021), this negative gradient means that wind streamlines are decreasing in velocity but increasing in height with increasing disc radius. As described in Section 5.3.2 this study measures the same negative gradient in all three years. As this measurement is made in the brown part of the position spectrum (Fig. 5.7) the question remains as to whether the gradient is a property of the LVC-M, LVC-L or the two components combined. Chou et al. (2025) associate this signal with the LVC-M for the following reasons. Firstly, the $[\text{S II}] \lambda 6731$ line has a strong negative gradient but does not have a strong contribution from the LVC-L. Secondly, the negative gradient is also strong in the 2003 $[\text{O I}] \lambda 6300$ line where the contribution from the LVC-L is again marginal. Thirdly, they point out that it could not be tracing one of the higher velocity components e.g. the LVC-H, as the negative gradient is seen in $[\text{O I}] \lambda 5577$ which does not trace these components. Overall they argue that the LVC-M emission contributes a larger velocity range within the negative velocity gradient in the position spectrum.

The analysis of the 2010 and 2021 LVC confirms all of the above points made by Chou et al. (2025). Even when a stronger contribution from the LVC-L begins to be seen in 2010 and 2021, the gradient of the position spectrum does not change. Therefore, it is also concluded here that the negative velocity gradient is a property of the LVC-M and that the LVC-M traces a disc wind. An important result from this work is that the velocity of this wind is stable compared to the velocity of the jet. It appears that the only effect the LVC-L has on the position spectrum is to reduce the size of the centroids in line with it originating from the upper disc atmosphere.

The next question to be addressed is whether the LVC-M traces a MHD disc wind or a photoevaporative wind. By comparing the line profiles and spatial scales of the LVC-M sub-component with the photoevaporative and MHD disc wind models of Weber et al. (2020), Chou et al. (2025) conclude that neither model can fully explain the LVC-M. In Whelan et al. (2021) the minimum de-projected vertical height of the LVC sub-component tracing the disc wind, is measured to be 2.7 AU. As this is well

within the gravitational potential well of the star, they argue that this rules out a photoevaporative wind component in this case. In Fig. 5.7, the LVC-M emission in 2003 [O I] $\lambda 5577$ is also traced to a minimum de-projected vertical height of 2 AU above the disc mid-plane in [O I] $\lambda 5577$, meaning that the same argument can be used here for the LVC-M. Interestingly, the lack of change in the velocity of the LVC could be taken as evidence that it is not linked to the jet as suggested by current MHD jet launching models.

5.5 Conclusions and Summary

The following are the main conclusions from this study

1. Up to six blueshifted components in the FEL line profiles alongside a red-shifted wing are identified in agreement with the study of Chou et al. (2025). The slowing of the DG Tau jet is confirmed and a decrease in velocity of $\sim 100 \text{ km s}^{-1}$ is seen from 2003 to 2021.
2. Further evidence that the HVC2 is a shock at the base of the DG Tau jet is presented. This includes a position for the HVC2 in agreement with the known location of the X-ray knot and a decrease in the velocity of the HVC2 from 2003 to 2010 of the same size as the decrease in the jet velocity.
3. The analysis of the LVC-H while limited by the spectral resolution of the data supports the idea of Chou et al. (2025) that it is material entrained by the jet.
4. The LVC-M and LVC-L are found to be variable in brightness and their peak velocity. However, the velocity gradient for these components does not change significantly between the years or the three lines.
5. In agreement with Chou et al. (2025) the negative velocity gradient recorded for the region of the LVC-M and LVC-L is associated with the LVC-M and it is concluded that this LVC sub-component is tracing a disc wind. The minimum de-projected height of LVC-M in [O I] $\lambda 5577$ of 2 AU favours an MHD wind.

6. It is particularly noteworthy that the velocity of the LVC-M does not change as the velocity of jet decreases.

Overall this study shows the potential of combining kinematic fitting and spectroastrometry to track variability and examine the origin of LVC sub-components. However, there are some key caveats. For a complex system like DG Tau there are limitations to the kinematic fitting due to blending between velocity components. The low time cadence of this data also limits the ability to assess if there is a time lag between the changes observed in the different velocity components in the outflow. However, these results do provide some very interesting new constraints for models of outflows emitted from YSOs. The results of this work especially highlight the need for future observational studies to investigate these outflows at a much higher cadence sampling in time (e.g. observations every few months instead of years). In particular this would be used to explore the possibility of a time lag between changes in the high velocity jet and changes in the low velocity wind in an MHD disc wind scenario.

Chapter 6

Interferometric analysis at the heart of IRAM 04191+1522: Uncovering complex molecular morphology & kinematics¹

6.1 Chapter Introduction

The dominant formation mechanism of BDs remains uncertain. The current literature divides the possible formation processes of BDs into two main branches: star-like or planet-like (see Chapter 1) In the planet-like scenario, the brown dwarf (BD) forms via turbulent fragmentation of material within a circumstellar disc surrounding a young protostar. This is followed by the subsequent ejection of the BD from the circumstellar disc via multiple body-body interactions (Reipurth and Mikkola, 2015). In the star-like scenario, BDs form via the gravitational collapse of fragmented dense cores of gas and dust within molecular clouds. Turbulence in these cores allows for the production of extreme density fluctuations within the core, decreasing the typical Jeans mass needed for the production of BDs within MCs (Padoan and Nordlund, 2002, 2004). Therefore, in order to determine the main formation mechanism of BDs and examine the role of outflows in the sub-stellar formation process, dedicated studies of the youngest and most embedded objects have to be conducted. The focus of this chapter is on one such young and heavily embedded object, IRAM 04191+1522. The objective of this analysis is to detect the presense of outflowing/infalling material moving out of/towards IRAM 04191+1522 in order to examine the role of outflows during the formation process of this object and to determine if this objects formation "pathway" is star-like or

¹Paper submitted to A&A and is in review.

planet-like.

The target, IRAM 04191+1522 (hereafter referred to as IRAM04191) is a young ($\sim 10^4$ years), highly embedded Class 0 system located in Taurus (distance ~ 140 pc) with an estimated protostellar mass of $0.05 M_{\odot}$ surrounded by an envelope of mass $0.5 M_{\text{sun}}$ (see Table 6.1 and references therein). It was discovered in 1999 via the detection of a large scale CO outflow (0.1 pc) in the southern part of the Taurus star forming region using the IRAM 30m telescope during a survey searching for embedded YSOs (see Fig. 6.1 and André, Motte, and Bacmann, 1999).

This is particularly interesting as the length scale of this molecular outflow is somewhat comparable to the parsec scale molecular outflows typically driven by Class 0 protostars (e.g. Eislöffel, 2000). Since its discovery, IRAM04191 has become a well studied object over the last two decades. Key findings from some of the previous literature surrounding IRAM04191 are as follows. After the discovery of IRAM04191, a study of the structure and kinematics of the envelope surrounding this source was conducted by Belloche et al. (2002). They find evidence of simultaneous infall, outflow and a differential rotation signature between the inner and outer envelope in several molecular species.

The first dedicated molecular line study concerning IRAM04191 was conducted by Lee, Ho, and White (2005) and Lee et al. (2002) and mapped the region surrounding IRAM04191 in the CO (1-0), HCO^+ (1-0) and N_2H^+ emission using the Berkeley-Illinois-Maryland Association Array (BIMA) at Hat Creek California. Lee, Ho, and White (2005) find that the CO (1-0) emission traces two structures along the outflow axis (axis A1 in their Fig. 1, position angle (PA) = 30°)².

1. A bipolar outflow along the north-east (red lobe) south-west (blue lobe) direction.

²Measured east of north.

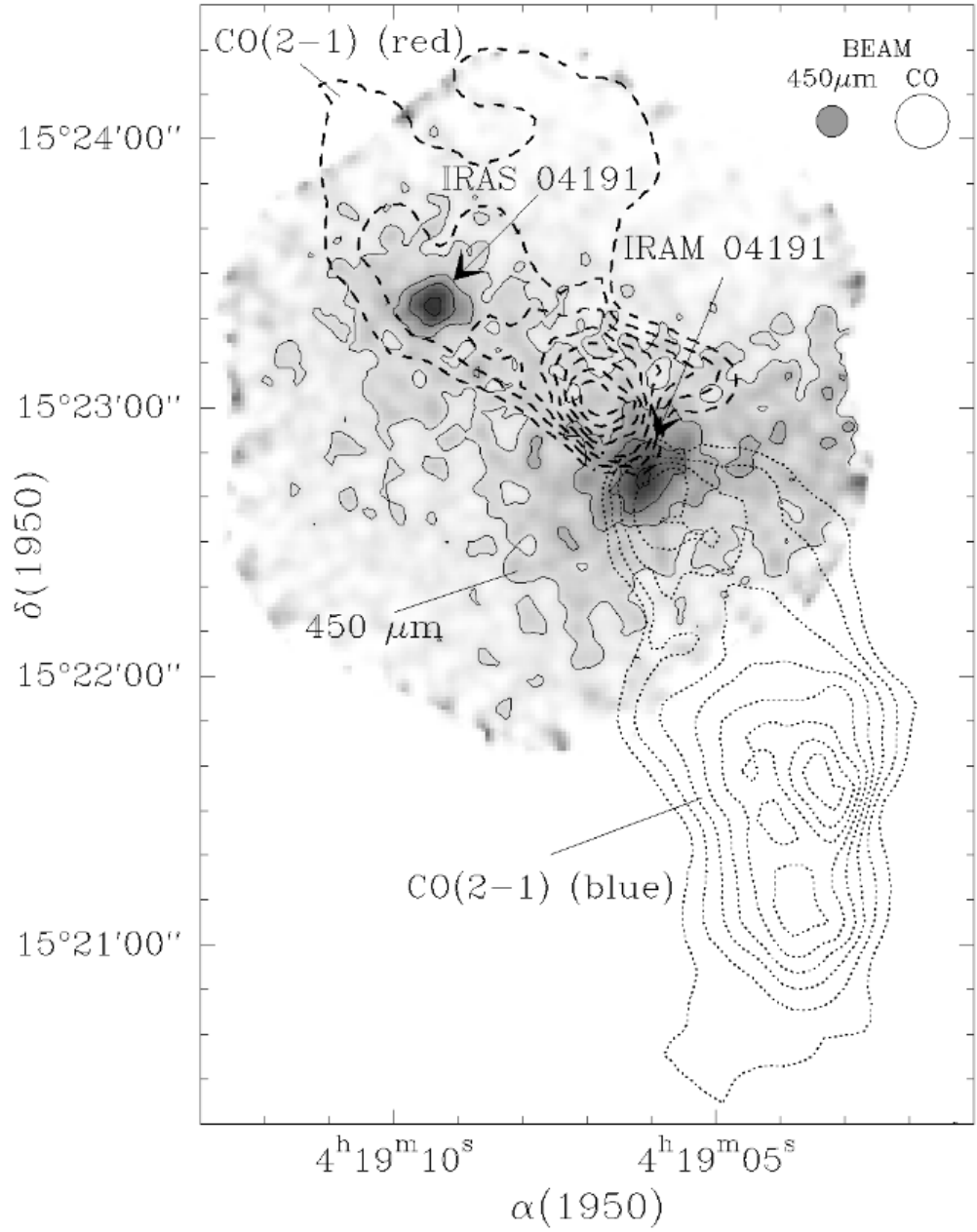


FIGURE 6.1: Taken from André, Motte, and Bacmann (1999). This is an IRAM 30m map of the CO (2–1) bipolar flow emanating from IRAM 04191 overlaid on a SCUBA 450 mm continuum image (gray scale). The blue-shifted CO (2–1) emission (dotted contours) is integrated over the 0 to 5 km s^{−1} velocity range, and the red-shifted CO (2–1) emission (dashed contours) is integrated over the 8 to 13 km s^{−1} velocity range. The CO (2–1) contours go from 5 to 21 K km s^{−1} in 2 K km s^{−1} intervals. The 450 μm solid contours go from 0.35 to 1.25 Jy per 7'' beam by 0.30 Jy beam^{−1}. The rms noise level is ∼ 0.09 Jy beam^{−1}. The peak flux densities of IRAM 04191 and IRAS 04191 are both 1.3 Jy per 7'' beam at λ = 450 μm.

2. Strong emission $\sim 20''$ (2800 au) to the north-east of the IRAM04191 source is detected, which they suggest is associated to an internal bow shock driven by an episodic collimated jet.

Lee, Ho, and White (2005) also draw a second axis (axis A2 in their Fig. 1, PA = 18°) from the cavity tip to the centre of the molecular outflow. This figure has been reproduced and is presented in Fig. 6.2. The authors suggest that this morphology may hint toward a binary nature for IRAM04191. A dense condensation of material in HCO^+ was also detected (labelled B2 in their Fig. 1) at a distance $40''$ (5600 au) to the south-west of IRAM04191. The authors state that this is likely linked to an unresolved expanding shell of gas interacting with the CO outflow. This condensation has also been detected in CS and in CH_3OH (Takakuwa, Ohashi, and Hirano, 2003). The HCO^+ emission was also found to trace a structure of $\sim 50''$ (7000 au) in length at a low velocity of -1 km s^{-1} to the south-west of IRAM04191 which may be produced via an interaction with the molecular jet. The detected N_2H^+ emission was attributed to clumpy dense core of material with a mean radius of $10''$ (1400 au) perpendicular to a molecular outflow traced by the CO emission.

Dunham et al. (2006) reported the first Spitzer observations of IRAM04191 using the InfraRed Array Camera (IRAC). They use these observations and radiative transfer models to show that IRAM04191 has an internal luminosity (i.e. excluding any luminosity arising from heating caused by an external radiation field) of $L_{\text{int}} = 0.08 \pm 0.04 L_\odot$. This is consistent with a Very Low Luminosity Object (VeLLO) and as such, this source is classified as a VeLLO (di Francesco et al., 2007; Vorobyov et al., 2017). The study of VeLLO driven outflows has been of particular interest in recent years (Kim et al., 2019) as a means of identifying sub-stellar mass candidates and studying their formation at the earliest stages (Santamaría-Miranda et al., 2021).

However, Dunham et al. (2006) also show that when placing the outflow parameters (e.g. bolometric luminosity, outflow mass and outflow force) derived in André, Motte, and Bacmann (1999) and Lee et al. (2002) on the plots relating these quantities (e.g.

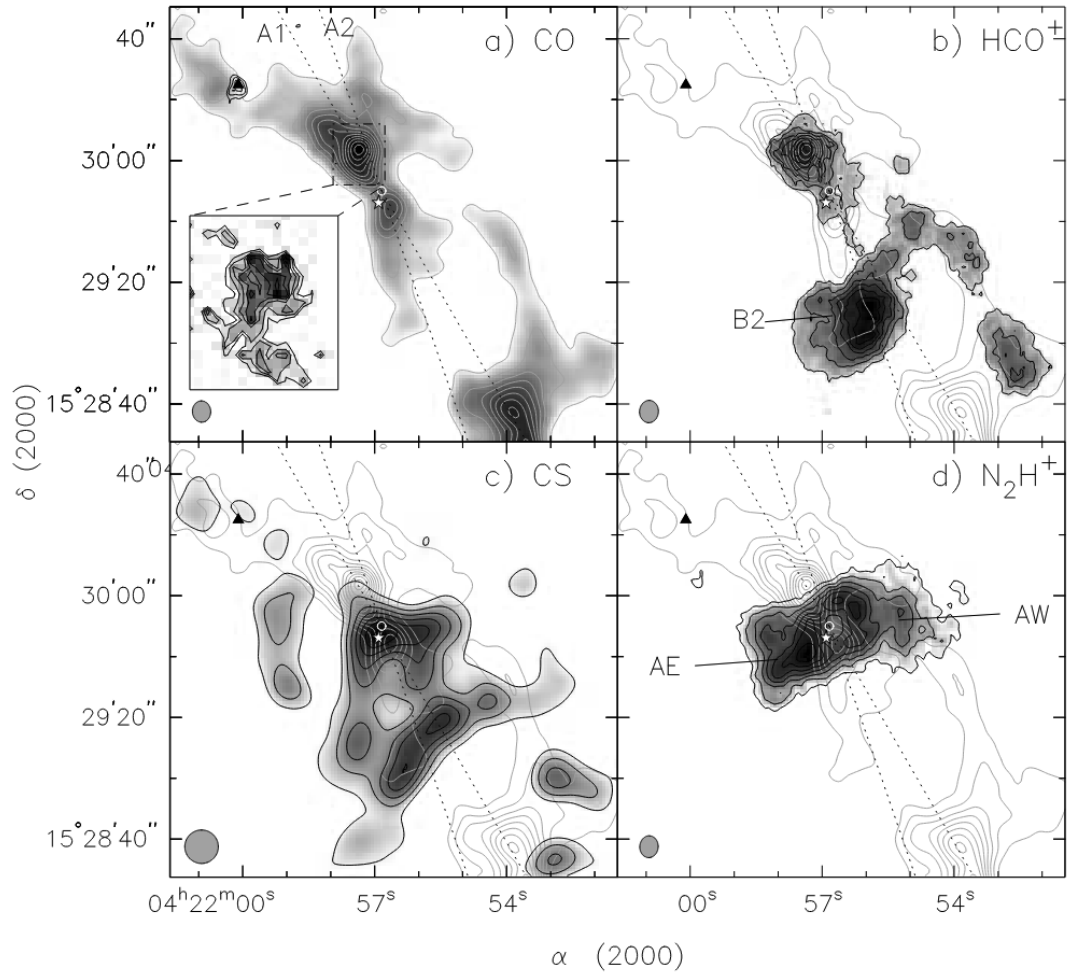


FIGURE 6.2: Figure 1 of Lee, Ho, and White (2005). Each panel shows the continuum (dust) emission as black contours. In the HCO^+ , CS and N_2H^+ panels the CO emission is marked in gray contours. The star indicates position of IRAM04191 as detected in 227 GHz continuum by the PdBI. The white circle shows the centre of the CO outflow. The triangle indicates the position of the Class I source IRAS 04191+1523. The dotted lines show two axes, A1 and A2, that connect the centre of the CO outflow to the tips of the two CO cavities (see Lee et al., 2002). The details regarding the contour levels and sigma clipping of these maps and the inset map are described in the original source.

Fig. 5 and 7 in Wu et al., 2004) they find that a source with an internal luminosity of $L_{\text{int}} \geq 1 L_{\odot}$ is needed to drive the large CO outflow driven by IRAM04191. This is inconsistent with the internal luminosity derived via their radiative transfer modelling. To explain this inconsistency, they hypothesise that their Spitzer/IRAC observations are a snapshot of IRAM04191 in a quiescent state and that this source may be subject to periods of episodic accretion.

The possible episodic mass accretion in IRAM04191 was further explored in Anderl et al. (2020) as part of the CALYPSO (Continuum And Lines in Young ProtoStellar Objects) survey using high angular resolution Plateau de Bure Interferometer (PdBI) observations of N_2H^+ and C^{18}O . Time dependant chemical and radiative transfer models were used to reproduce the line profiles extracted through the emission peaks observed in the integrated intensity maps of the N_2H^+ and C^{18}O emission lines (see their Fig. 1). Their results indicate that these line profiles cannot be reproduced using a constant luminosity model based on the current value of internal luminosity for IRAM04191 ($0.08 \pm 0.04 L_\odot$). Instead, the peak in the N_2H^+ emission at $10''$ from the source may be caused by a past accretion burst event producing a luminosity 150 times greater than the present day value. Their models predict that if such accretion bursts occur for $\sim 1\%$ of the Class 0 lifetime of IRAM04191 (assumed to be $0.4 - 0.5 \times 10^6$ years; Evans et al., 2009) then, the mass of IRAM04191 will reach $0.2 - 0.25 M_\odot$ by the end of the Class 0 stage, which is above the sub-stellar mass limit.

Lastly, the work of Podio et al. (2021) gives more insight into this source. They performed high angular resolution ($0''.5$ to $1''$) CALYPSO PdBI observations of IRAM04191 in CO (2–1), SO 6_5-5_4 , and SiO (5–4). A redshifted outflow is detected only in CO (2–1) from $V_{lsr} = 8.2 \text{ km s}^{-1}$ to 21.2 km s^{-1} , extending $\sim 4''$ to the north-east at a PA of 20° .

In summary, the literature surrounding IRAM04191 has revealed a vast amount of information about this source. IRAM04191 is surrounded by a large, differentially rotating envelope and the central source shows hints of possible episodic mass accretion. There is also evidence of a large 0.1 pc outflow in CO. The combination of an outflow and lots of surrounding material available for accretion onto IRAM04191 suggests that IRAM04191 may be able to gather enough surrounding material to evolve into a low mass YSO instead of a brown dwarf by the end of the Class 0 stage (Anderl et al., 2020). Hence, IRAM04191 may be a cornerstone object in understanding the link between the formation processes of sub-stellar and stellar mass objects.

The aim of this chapter is to provide a detailed study of the molecular gas emission in close proximity to IRAM04191 at an angular resolution ($0''.27$) better than what is available in the current literature. A dedicated analysis of the molecular emission lines tracing outflowing and infalling gas in close proximity to this source may hold critical evidence needed to study a possible formation pathway for IRAM04191 and examine if its formation process is star-like or planet-like.

This study focuses on the morphology and kinematics of three molecular emission lines ^{13}CO (3–2), C^{18}O (2–1) and SO 6_5-5_4 (Table 6.2), which are known to trace outflowing and/or infalling gas in young protostellar systems (Garufi et al., 2022; Hales et al., 2024; Podio et al., 2021; Tychoniec et al., 2021). These lines were chosen based on their ability to trace cold/dense gas in close proximity of IRAM04191 (from a few 10s to several 100s of AU). The transition selected for each line was chosen based on the available public archival 12m array data for IRAM04191. The following paragraphs provide a brief description of what each molecular line presented in this chapter can trace according to current literature.

The ^{13}CO (3–2) emission line traces dense ($\sim \times 10^3 - 10^6 \text{ cm}^{-3}$) and warm (few 10s to 100 K) gas. ^{13}CO (3–2) is an isotopologue of the more abundant CO molecule and as such, it traces denser gas closer to the emitting source than CO. The literature has shown that transitions of ^{13}CO emission have been found in molecular outflows emanating from very low mass stars (Santamaría-Miranda et al., 2020) and in thin filamentary structures within molecular clouds (Neupane et al., 2024).

C^{18}O (2–1) is another isotopologue of CO and a tracer of gas at temperatures of ~ 15 K. C^{18}O (2–1) is less abundant than CO and ^{13}CO hence, it traces even denser gas closer to the emitting source. The 2–1 transition of C^{18}O has an upper level energy of ~ 15 K (Table 6.2) while the 3–2 transition of C^{18}O has an upper level energy of ~ 32 K. This makes our selected transition of C^{18}O particularly useful for detecting cold gas in close proximity to our target. The literature has shown that transitions of C^{18}O gas

TABLE 6.1: Properties of IRAM 04191+1522.

Property	Value	Reference
Source mass ($M_{\odot}$)	0.05	Anderl et al. (2020)
Env. Mass ($M_{\odot}$)	0.5	Maury et al. (2019)
Age (yrs.)	$0.8 - 2 (\times 10^4)$	André, Motte, and Bacmann (1999) ^a
$\dot{M}_{\text{acc}}$ ($M_{\odot}/\text{yr}$)	$2 \pm 1 (\times 10^{-6})$	Kim et al. (2019)
Disc incl. ($^{\circ}$)	50	Belloche et al. (2002)
Distance (pc)	140	André, Motte, and Bacmann (1999)
V_{sys} (km s^{-1})	6.5 ± 0.4	this work ^b
L_{int} ($L_{\odot}$)	0.08 ± 0.04	Dunham et al. (2006)

^a = The authors estimate the lower limit for the age of IRAM 04191+1522 based on the dynamical timescale of the outflow whereas, the upper limit for the age was taken from the evolutionary tracks of Saraceno et al. (1996) and Myers et al. (1998).

^b = The systemic velocity was estimated from the line spectra presented in this work. The uncertainty is taken as the spectral resolution of the ^{13}CO (3–2) cube as it has the lowest spectral resolution of the ALMA data that is presented (see Table 6.4).

emission have been observed in several different structures around low mass YSOs and BDs. It has been found in the cold envelope regions surrounding young stellar systems (Tychoniec et al., 2021) and in the discs surrounding young protostars and BDs (Alves et al., 2017; Ruaud, Gorti, and Hollenbach, 2022).

The SO 6₅–5₄ molecular emission line has been detected in shocked regions, close to the circumstellar disc (Dutrey et al., 2024). Emission from oxygen bearing species such as SO have also been observed in the knots within the high velocity molecular jet launched by low mass young stars and in the jet bow shocks (Frank et al., 2014; Lee et al., 2010; Tychoniec et al., 2021). This molecule is also associated to streamers, as either infalling rotating structures (Flores et al., 2023) or in shocks where the streamer impacts flattened disc like structures (Kido et al., 2023).

This chapter is organised as follows: Section 6.2 describes the observations and the steps used to reduce and analyse these data. In section 6.3 presents the results from the analysis of the data, including the morphological and kinematic properties of the reported emission lines. Section 6.4 provides a discussion of the results. Lastly, section 6.5 summerises the findings and draws conclusions from them.

TABLE 6.2: Properties of the emission lines used in this study.
Taken from the Advanced Splatalogue database available at <https://splatalogue.online/#/home>.

Species	Transition	Rest Frequency [GHz]	E_{low} [K]	E_{up} [K]
^{13}CO	3–2	330.588	15.9	31.7
C^{18}O	2–1	219.560	5.3	15.8
SO	6 ₅ –5 ₄	219.949	24.4	35.0

6.2 Observations, data reduction & analysis

IRAM04191 was observed in the millimetre and sub-millimetre range (Band 6 and 7) with ALMA in 2016 and in 2019 (Table 6.3). The molecular line emission data presented here has been obtained from the ALMA science archive from two different observing programmes. The C^{18}O (2–1) and SO 6₅–5₄ emission lines were targeted in 2016 (ID: 2016.1.01284.S, PI: A. Maury) and the ^{13}CO (3–2) line was targeted in 2019 (ID: 2019.1.00847.S, PI: P. Sheehan). The calibrated measurement sets (CMs) for these data were obtained from the European ALMA ARC through the European ARC calibrated data service. The CMs for 2016 the 2019 data were generated using version 2024.1.0.8 and version 42866M (Pipeline-CASA56-P1-B) of the ALMA pipeline respectively. Following this, the CASA, (McMullin et al. (2007), version 6.5.5) package was used to produce the image cubes for each emission line via the `tclean` task. A Briggs weighting scheme and robust factors of 0.5, 1.0 and 2.0 were chosen for imaging the C^{18}O (2–1), SO 6₅–5₄ and ^{13}CO (3–2) lines respectively. These robust factors were selected in order to recover as much flux as possible from the gas surrounding the central source while preserving the spatial resolution required to achieve the scientific objectives. The image cubes were not primary beam corrected. A summary of the final image cube parameters is provided in Table 6.4.

After the production of the ^{13}CO (3–2) line emission image cube, it was noted that the phase centre (RA = 4^h21^m56.40^s, Dec = 15°29′48.90″ in J2000) of the image cube was not centred at the position of the central source. This was corrected using the `phaseshift` task in CASA, to the position of IRAM04191 (RA = 4^h21^m56.90^s, Dec =

15°29^m46.05^s in J2000) so, that it matched the phase centres of the C¹⁸O (2–1) and SO 6₅–5₄ image cubes. The channel maps, moment maps and PV diagrams presented in the results sections were produced from these final emission line cubes using a combination of tools available in CARTA (version 4.1; Comrie et al., 2021; Wang et al., 2026) and custom *Python* scripts. The channel maps for each of the three molecules presented are in Appendix B.

In addition to the ALMA data presented, these data are complimented with the archival Spitzer observations originally presented in Dunham et al. (2006)³. The details of the data reduction, analysis, and interpretation of these IRAC observations are presented in Dunham et al. (2006). These data have been added to this chapter to aid in the interpretation and discussion of the ALMA archival data presented here.

³The Spitzer space telescope archival data is available at https://irsa.ipac.caltech.edu/applications/Spitzer/SHA/?__action=layout.showDropDown&

TABLE 6.3: Summary of ALMA observations.

Array	Date [yyyy-mm-dd]	Project ID	PWV [mm]	Flux Cal.	Phase Cal.	Bandpass Cal.	Avg. Elevation [°]	Intr. Time [s]	AR [arcsec]	MRS [arcsec]	Baseline (min-max) [m]
12m	2016-10-15	2016.1.01284.S	0.641	J0510+1800	J0433+0521	J0425+1755	37.5	3991.680	0.154	1.949	18 - 1809
12m	2019-10-07	2019.1.00847.S	0.436	J0510+1800	J0431+1731	J0431+2037	46.7	907.200	0.285	3.758	15 - 1547

AR = Angular Resolution. AR is defined as the smallest scale structure that can be resolved by the interferometer. This is determined by the length of the longest baseline and the wavelength of the observation.

MRS = Maximum Recoverable Scale. MRS is defined as the largest scale structure that can be resolved by the interferometer. This is determined by the length of the shortest baseline and the wavelength of the observation.

TABLE 6.4: Properties of the produced emission line cubes.

Species	Transition	Robust	RA Range [J2000]	Dec Range [J2000]	Beam Major Axis [arcsec]	Beam Minor Axis [arcsec]	Beam PA [°]	RMS [mJy/beam]	Spectral Res. [km s ⁻¹]
¹³ CO	3-2	2.0	04:21:56.076, 04:21:57.734	+15.29.34.049, +15.29.58.009	0.39	0.33	155.6	9.484	0.442
C ¹⁸ O	2-1	0.5	04:21:56.739, 04:21:57.022	+15.29.43.952, +15.29.48.044	0.19	0.17	102.1	2.623	0.166
SO	6 ₅ -5 ₄	1.0	04:21:56.548, 04:21:57.255	+15.29.40.990, +15.29.51.210	0.24	0.20	119.4	2.102	0.166

RA Range = Right Ascension extent, Dec Range = Declination extent, PA = Position Angle (measured east of north), RMS = Root mean square noise, Res. = Resolution

TABLE 6.5: Fitting results for the emission lines of interest done in CARTA. The uncertainty range on each measurement are in parentheses.

Line	Comp.	V_c [km s ⁻¹]	Intensity [mJy]	FWHM [km s ⁻¹]	Intgr. Flux [mJy × km s ⁻¹]
¹³ CO (3-2)	1	5.63 (0.05)	67.89 (6.71)	0.87 (0.11)	63.19 (8.05)
	2	7.66 (0.09)	56.40 (4.44)	2.02 (0.23)	121.04 (11.05)
C ¹⁸ O (2-1)	1	5.66 (0.04)	8.29 (0.42)	1.07 (0.08)	9.44 (0.73)
	2	7.45 (0.07)	5.33 (0.34)	1.66 (0.18)	9.43 (0.89)
SO 6 ₅ – 5 ₄	1	5.75 (0.11)	3.99 (0.45)	1.30 (0.25)	5.55 (1.03)
	2	7.48 (0.15)	2.93 (0.45)	1.34 (0.35)	4.16 (1.04)

6.3 Results

To preface the results section, it is noted here that the systemic velocity adopted for this chapter is 6.5 km s⁻¹ (Table 6.1). This was estimated from the spectra of each emission line which are presented later in the chapter.

6.3.1 ¹³CO (3–2) Morphology

In Fig. 6.3 (top row) presents the integrated and maximum intensity moment maps for the blueshifted emission traced by ¹³CO (3–2). In both maps two main structures traced by the blueshifted ¹³CO (3–2) gas are observed. The first is a large ring shaped filament like structure that extends $\sim 6''$ (measured from the 1σ contour) to the north-west. The second is smaller oblong shaped structure centred at the position of the source that extends at most $1''$ along the north-south direction. In the blueshifted ¹³CO (3–2) maximum intensity map there is a hint that this oblong structure may be connected to nearby filaments of material.

Figure 6.4 (top row) shows the integrated and maximum intensity maps for the redshifted ¹³CO (3–2) gas. The morphology of the redshifted gas is strikingly different compared to the blueshifted gas. The redshifted side of this emission line shows several structures.

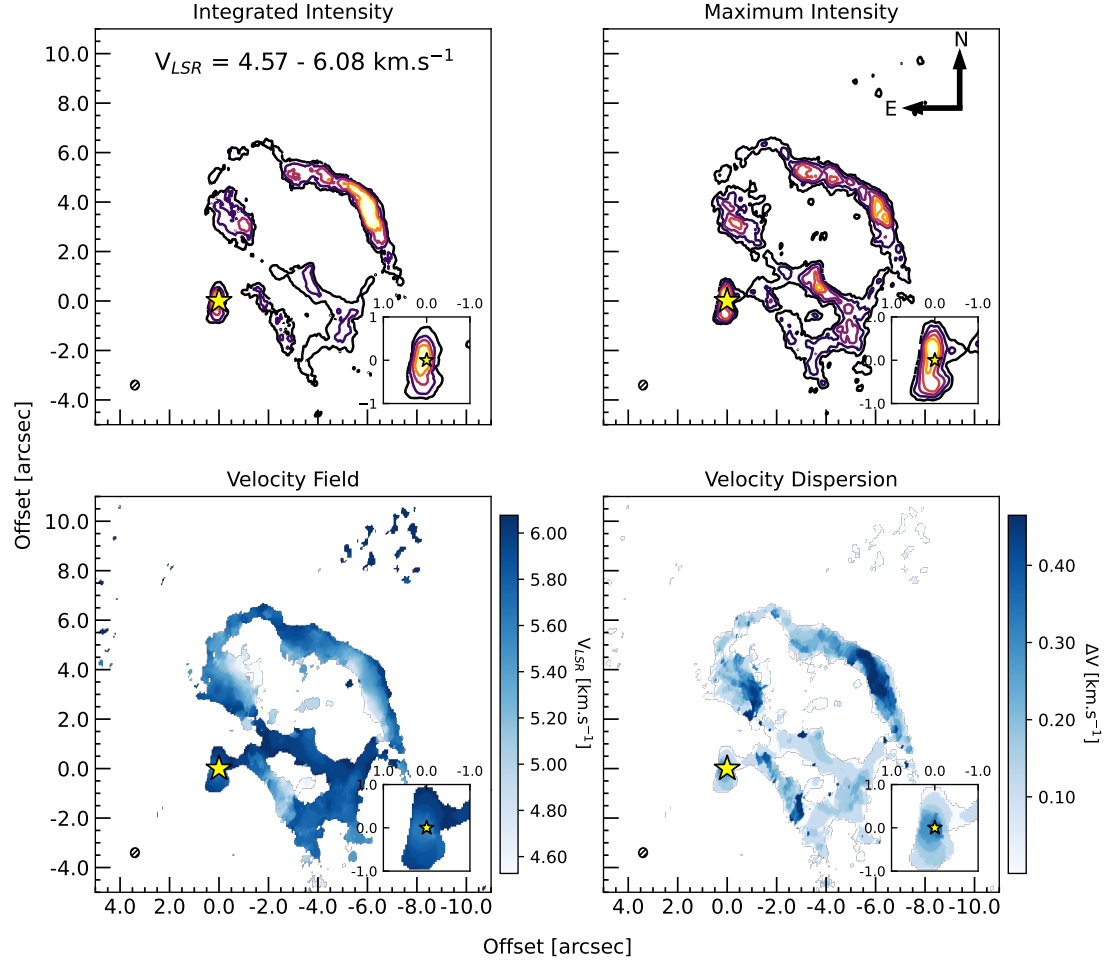


FIGURE 6.3: ^{13}CO (3-2) blueshifted emission moment maps integrated from 4.57 to 6.08 km s^{-1} . **Top left:** Integrated intensity map. Contours used were 1, 2, 3, 4 and 5σ ($\sigma = 14.38 \text{ mJy/beam} \times \text{km s}^{-1}$). **Top right:** Maximum intensity map. Contours used were 2, 3, 4, 5, 6 and 7σ ($\sigma = 13.2 \text{ mJy/beam}$). **Bottom left:** Velocity field map. **Bottom right:** Velocity dispersion map. In each map there is a small inset plot showing the innermost 2'' centred on the source. The source position is denoted by a yellow star. The beam is represented as a black ellipse in the bottom left corner of each map ($0''.39 \times 0''.33$ at 155.6°). The systemic velocity is 6.5 km s^{-1} (Table 6.1).

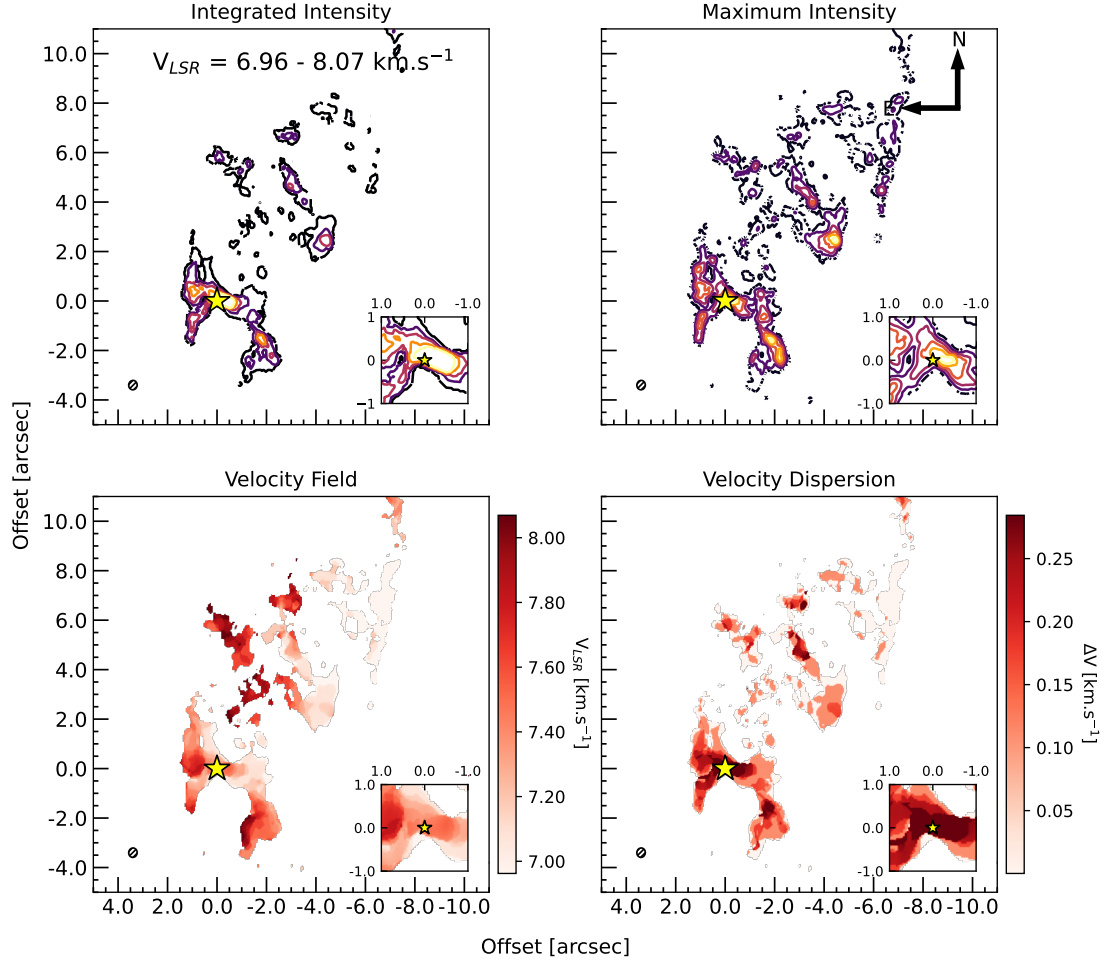


FIGURE 6.4: ^{13}CO (3–2) redshifted emission moment maps integrated from 6.96 to 8.07 km s^{-1} . **Top left:** Integrated intensity map. Contours used were 1, 2, 3, 4 and 5σ ($\sigma = 12.58 \text{ mJy/beam} \times \text{km s}^{-1}$). **Top right:** Maximum intensity map. Contours used were 3, 4, 5, 6, 7 and 8σ ($\sigma = 10.7 \text{ mJy/beam}$). **Bottom left:** Velocity field map. **Bottom right:** Velocity dispersion map. In each map there is a small inset plot showing the innermost $2''$ centred on the source. The source position is denoted by a yellow star. The beam is represented as a black ellipse in the bottom left corner of each map ($0''.39 \times 0''.33$ at 155.6°). The systemic velocity is 6.5 km s^{-1} (Table 6.1).

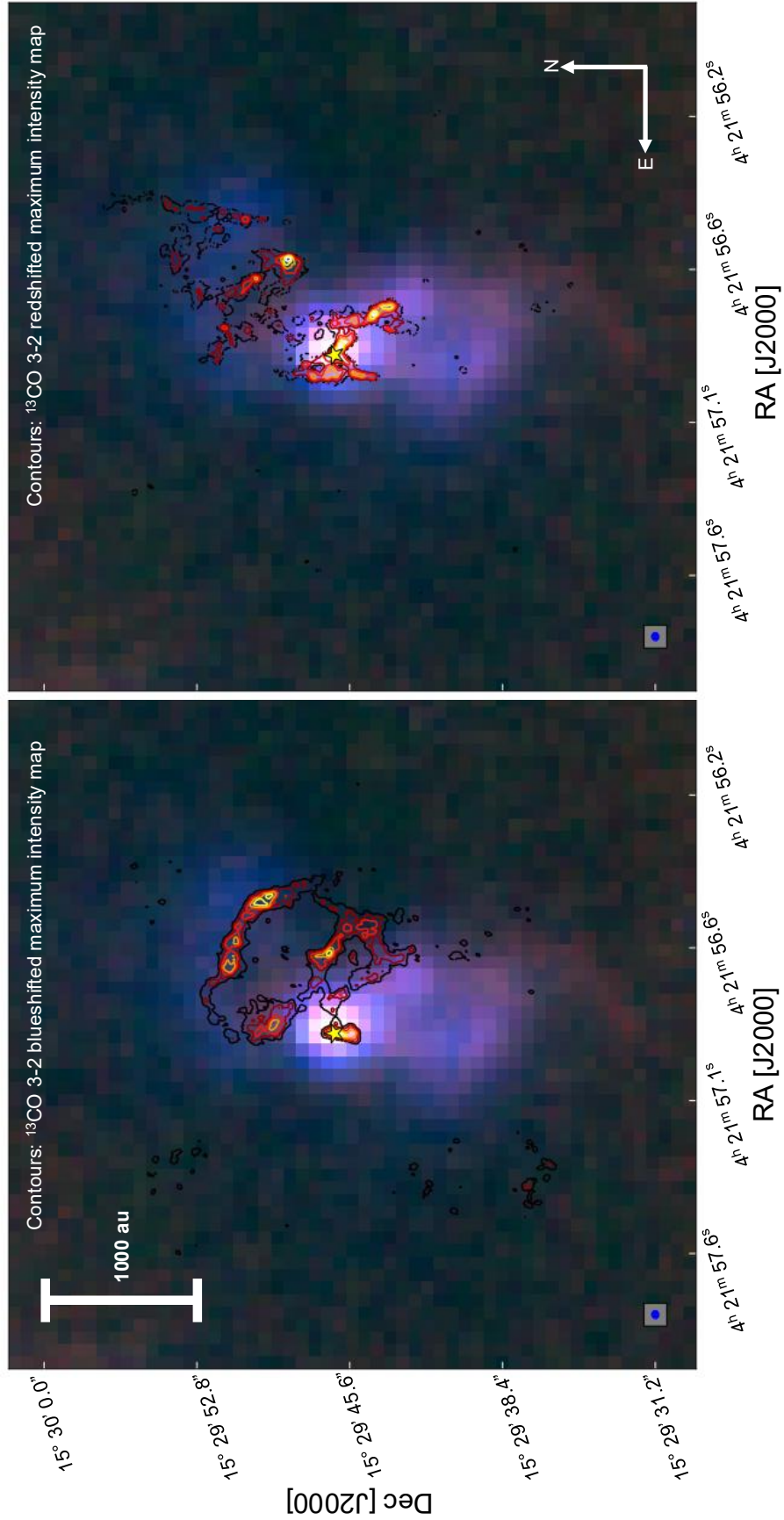


FIGURE 6.5: **Left:** Spitzer/IRAC RGB image consisting of the 4.5 (blue), 5.8 (green) and 8 (red) μm wideband filters compared to the blueshifted maximum intensity emission map of ^{13}CO (3–2) integrated from $V_{\text{LSR}} = 4.57$ to 6.08 km s^{-1} . Contours used were 2, 3, 4, 5, 6 and 7σ ($\sigma = 13.2 \text{ mJy}$). **Right:** The same IRAC RGB image compared to the redshifted maximum intensity emission map of ^{13}CO (3–2) integrated from $V_{\text{LSR}} = 6.96$ to 8.07 km s^{-1} . Contours used were 3, 4, 5, 6, 7 and 8σ ($\sigma = 10.7 \text{ mJy}$). The beam size ($0''.39 \times 0''.33$ at 155.6°) is shown by the blue ellipse in the bottom left corner of each map. The yellow star marks the source position. The systemic velocity for IIRAM04191 is 6.5 km s^{-1} (Table 6.1).

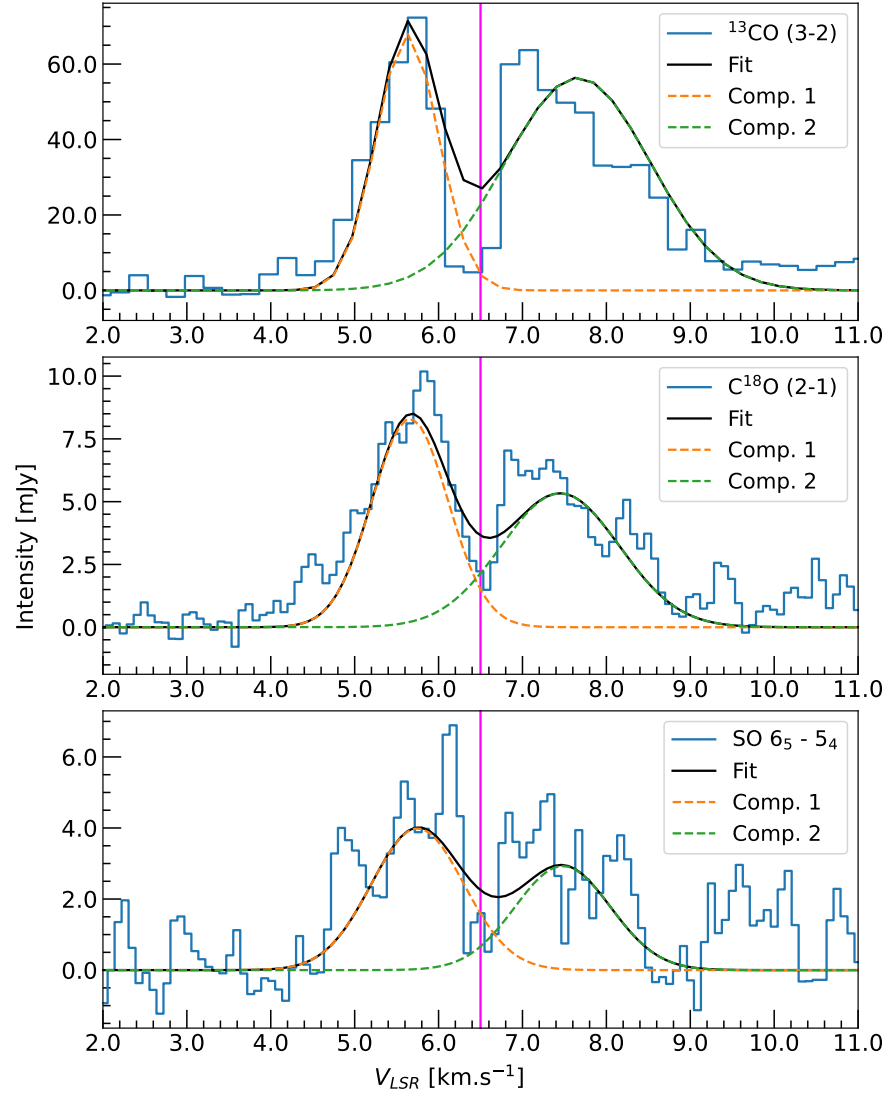


FIGURE 6.6: Fitting of the ^{13}CO (3-2) line (top) and C^{18}O (2-1) (middle) and $\text{SO } 6_5 - 5_4$ (bottom) lines done using CARTA after continuum subtraction. The spectra were extracted from a circular aperture of radius $0''.2$ centred at the source position. Each Gaussian component in the fit of each line is labelled as either 1 or 2. The magenta line corresponds to the systemic velocity (6.5 km s^{-1} , Table 6.1)

. The fitting results are presented in Table 6.5.

1. Sporadic emission is seen from $\sim 2''$ away from the source to around $10''$ in the north-west. This was measured from the $1\sigma/3\sigma$ contour of the integrated/ maximum intensity map of this emission.
2. Closer to the source, a straightened structure to the east of the source that extends along the north-south direction from $\sim -2''$ to $2''$ is seen, measured in the same way.
3. To the west of the source the redshifted emission begins to curve downwards to around $-3''.5$, measured in the same way as outlined above.

The structures observed to the east and west of IRAM04191 in the redshifted ^{13}CO (3–2) gas are clearly connected and may be part of the same structure. This is now referred to as the south-western structure. This larger structure in the south-west looks to be separate from the sporadic ^{13}CO (3–2) redshifted emission seen in the north-west.

Fig. 6.5 compares the blue and redshifted maximum intensity maps of ^{13}CO (3–2) to the Spitzer/IRAC observations originally presented and discussed in Dunham et al. (2006). In the left panel, the blueshifted ring shaped structure traced by ^{13}CO (3–2) appears to align well with the curved structure seen in the north-west of the IRAC image. This curved structure appears to be primarily traced by the $4.5\ \mu\text{m}$ filter. As noted in Dunham et al. (2006) the $4.5\ \mu\text{m}$ wideband contains three of the brightest H_2 emission lines (i.e. 0–0 S(11) at $4.18\ \mu\text{m}$, 0–0 S(10) at $4.40\ \mu\text{m}$ and 0–0 S(9) at $4.18\ \mu\text{m}$) and H_2 lines are known tracers of shocked material (Noriega-Crespo et al., 2014). This spatial overlap suggests that ^{13}CO (3–2) blueshifted ring shaped structure could be tracing shocked gas, as seen before in GMCs and filamentary structures (Zhou et al., 2023).

In the right panel of Fig. 6.5, the redshifted spurious ^{13}CO (3–2) emission appears to be spread all over the curved structure in the north-west traced in the $4.5\ \mu\text{m}$ filter of Spitzer/IRAC. The south-western structure in the red ^{13}CO (3–2) gas seems to tentatively overlap with the base of some extended emission in the south/south-west traced by the 4.5 and $8\ \mu\text{m}$ filters in the IRAC image. Also, as outlined in Dunham et

al. (2006) (their Fig. 1), the $4.5\ \mu\text{m}$ emission likely arises from an outflow cavity while the $8\ \mu\text{m}$ filter traces more diffuse emission due to polycyclic aromatic hydrocarbons (PAHs). Given this, the overlap between the south-western structure in the redshifted ^{13}CO (3–2) gas and the south/south-western extended emission in the IRAC image suggests that this redshifted ^{13}CO (3–2) gas may be tracing the base of an outflow cavity.

6.3.2 ^{13}CO (3–2) Kinematics

Figure 6.3 (bottom row) presents the velocity field and velocity dispersion maps for the blueshifted ^{13}CO (3–2) gas. These maps are rather complex but, the following is noted.

1. In the velocity field map (Fig. 6.3, bottom left panel) the ring shaped filamentary structure shows a complex velocity field throughout. The top half of the ring that faces north-west shows an inner edge that is at a lower local standard of rest (LSR) velocity (~ 4.6 to $5.4\ \text{km s}^{-1}$) compared to the outer edge of the ring (~ 5.6 to $6.0\ \text{km s}^{-1}$).
2. The inset plot at the bottom right of the velocity field map shows the innermost $2''$ closest to the source (i.e. the oblong shape). It was observed that V_{LSR} decreases from $\sim 6.0\ \text{km s}^{-1}$ to about $5.6\ \text{km s}^{-1}$ at the source position.
3. In the velocity dispersion map (Fig. 6.3, bottom right panel) the edge of the ring shaped filamentary structure in the north-west has the largest area of large velocity dispersion with values of $\Delta V \geq 0.40\ \text{km s}^{-1}$.
4. The inset plot at the bottom right of the velocity dispersion map shows the innermost $2''$ closest to the source (i.e. the oblong shape). An increase in velocity dispersion of the gas is observed. It increases from $\Delta V \sim 0.05\ \text{km s}^{-1}$ at $1''$ away from the source to $\Delta V \sim 0.35\ \text{km s}^{-1}$ at the source position. This increase in velocity dispersion suggests increased turbulence in the gas close to the source. This may be due to the launching of a jet/outflow or the impact of in-falling material onto a circumstellar disc (Thieme et al., 2022; Vazzano et al., 2021).

5. It is also noted here the innermost region of this inset map shows slightly different velocity dispersion values to the east and west of the source position, with the east side having slightly larger velocity dispersion.

The velocity field and velocity dispersion maps for the redshifted ^{13}CO (3–2) emission are shown in Fig. 6.4 (bottom row). As for the blueshifted emission, the velocity field and dispersion maps are complex however the following was noted.

1. In the velocity field map (Fig. 6.4, bottom left panel) the spurious emission to the north-west is mixed in terms of its velocity. The most extended emission to the north-west is at $\sim V_{lsr} = 7.0 \text{ km s}^{-1}$. This velocity increases as moving closer to the south-western structure that surrounds the source.
2. In the structure to the south-west, a complex velocity field is observed. The velocity field appears fastest at the eastern and south-western edge of this structure, but the field mixes into different velocities as moving closer to the source.
3. The inset map, showing the innermost 2" from the source, at the bottom right of this velocity field map shows that the velocity is higher to the east of the source ($\sim 7.8 \text{ km s}^{-1}$) than to the west of the source ($\sim 7.4 \text{ km s}^{-1}$).
4. In general, the velocity dispersion map for the redshifted ^{13}CO (3–2) emission shows the highest dispersion in the innermost 2" region nearest the source.

Lastly, Fig. 6.6 (top panel) presents the line profile extracted from the ^{13}CO (3–2) emission line cube from the innermost 0".2 from the source. The ^{13}CO (3–2) line profile is double peaked. This profile has been fitted with two Gaussian components (see Table 6.5). The blueshifted component of the fit is narrower and has a higher intensity than the redshifted component of the fit. The double peaked shape of this profile may indicate the presence of gas associated to a circumstellar disc in rotation or envelope material in rotation. The difference in shape of the blue and redshifted side of the line is interesting, as an extended red wing is seen, which may be indicative of an outflow (Ray and Ferreira, 2021). However, it is suspected that the shape of this line profile may be caused by a mix of different emission components in the ^{13}CO (3–2) gas in

the innermost region. This is supported by the complex morphology, velocity field and velocity dispersion maps for both the red and blueshifted ^{13}CO (3–2) gas in the 2'' region closest to the central source.

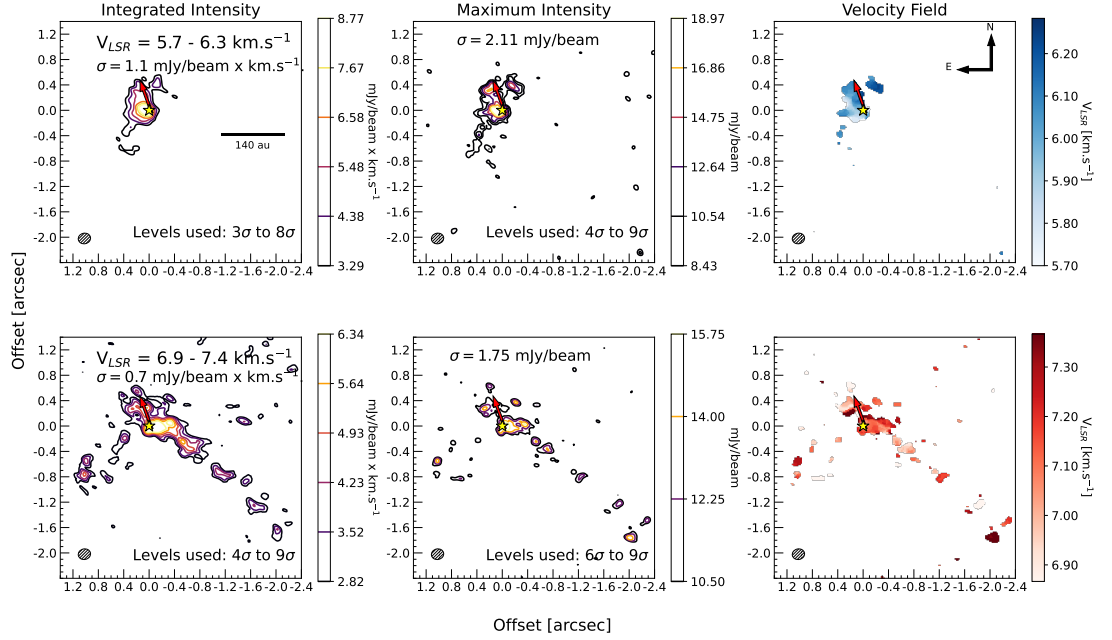


FIGURE 6.7: **Left column:** Integrated intensity (moment 0) maps for C^{18}O emission line. **Middle column:** maximum intensity (moment 8) maps for the C^{18}O emission line. **Right column:** Velocity field (moment 1) maps for the C^{18}O emission line. The velocity channels used for each row of maps is shown at the top of each map in the left column. The σ value for each moment 0 and 8 map is shown at the top of each respective map. The σ levels used for the contours in the moment 0 and 8 maps are in 1σ increments and the levels used are shown in the bottom right corner of each map. The beam ($0''.19 \times 0''.17$ at 102.1°) is represented as a black ellipse in the bottom right corner of every map. The source position is denoted as a yellow star. The red arrow shows the position angle (20°) of the redshifted outflow previously detected by Podio et al. (2021).

6.3.3 C^{18}O (2–1) Morphology

Figure 6.7 shows the integrated intensity (left column), maximum intensity (middle column) and velocity field (right column) maps over two velocity ranges (5.7 to 6.3 km s^{-1} and 6.9 to 7.4 km s^{-1}). These channels were selected as the most prominent C^{18}O (2–1) emission features are detected in the blue and redshifted side of the C^{18}O (2–1) emission line in these channels. This section discusses the results from the

integrated and maximum intensity maps over these two selected regions.

The morphology of the $5.7 - 6.3 \text{ km s}^{-1}$ velocity channels are discussed first. In both the integrated and maximum intensity maps, there are two extended components in the C^{18}O (2–1) gas. A north-eastern component was detected. It was observed out to $\sim 0''.5$. It is noted that this material lies close to the PA (30°) of the previously detected large scale outflow (André, Motte, and Bacmann, 1999; Kim et al., 2019; Podio et al., 2021). A second component is also observed along the north-west to south-east direction. The north-western part is of $\sim 0''.6$ while the south-eastern part of this component is observed out to $\sim 0''.7$. The total projected length of this component is $\sim 1''.3$. These distances are measured from the 3σ contour of the integrated intensity map.

The second velocity channel region selected for the C^{18}O (2–1) emission line (6.9 to 7.4 km s^{-1}) also shows two components and they are the most extended and collimated components observed in the C^{18}O (2–1) gas. The south-eastern component seen in the 5.7 to 6.3 km s^{-1} channels is also observed here in the 6.9 to 7.4 km s^{-1} velocity channels. The PA has changed slightly and this component now extends out to $\sim 0''.8$. A new component is observed in the south-west in the 6.9 to 7.4 km s^{-1} velocity channels. This additional south-western component is narrow knotty like emission seen out to $\sim 2''$ from the source in both the integrated and maximum intensity maps. These knot like features may appear due to the high angular resolution of the observations (emission more extended than $2''$ in this line seems to be filtered-out by the interferometer) and faintness of the emission.

From the analysis of the 6.9 to 7.4 km s^{-1} velocity channels in C^{18}O (2–1) there is a comparable morphology between the C^{18}O (2–1) gas in these channels and the red-shifted emission seen in ^{13}CO (3–2). In Fig. 6.8 the morphology of these two emissions are compared. A striking similarity in the shape of the gas traced in the south-east by

both lines is observed. The C^{18}O (2–1) emission in this velocity range appears to fit inside the redshifted ^{13}CO (3–2) emission. As previously mentioned in section 6.3.1, the redshifted ^{13}CO (3–2) gas may be tracing the base of an outflow cavity. The alignment of these two emissions suggests that the C^{18}O (2–1) in these velocity channels may be tracing a denser region of gas at the base of an outflow cavity. This is interesting given there is recent evidence of nested outflow morphology of atomic and molecular emission in several young low mass YSOs (Birney et al., 2024a; Pascucci et al., 2025). However, it is noted here that this effect may in part, be due to the fact that C^{18}O (2–1) traces colder and denser regions than ^{13}CO (3–2).

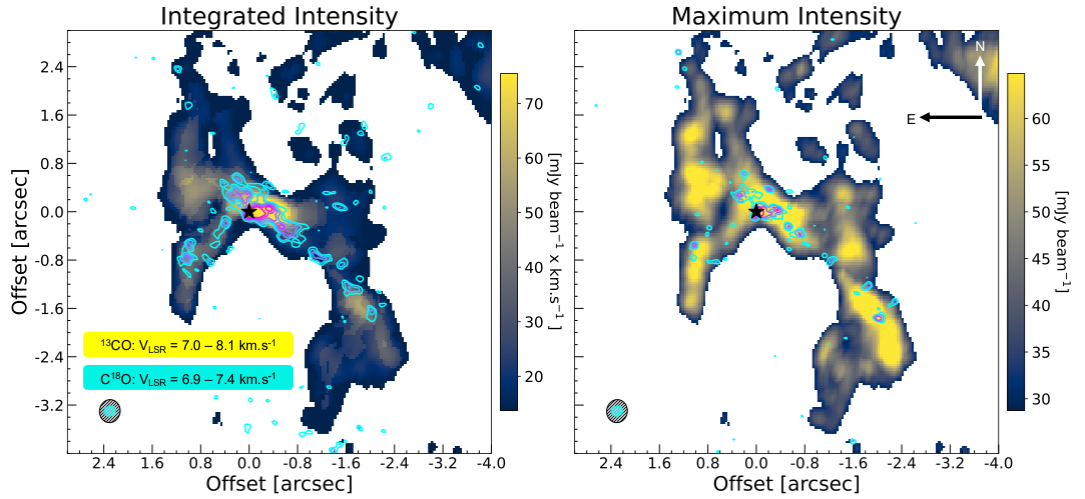


FIGURE 6.8: Comparison of the redshifted emission of the ^{13}CO (3–2) line from $V_{\text{LSR}} = 7.0 \text{ km s}^{-1}$ to 8.1 km s^{-1} (represented as a raster map) to the C^{18}O (2–1) line emission integrated from $V_{\text{LSR}} = 6.9$ to 7.4 km s^{-1} (represented as the cyan and purple contours). Left = Integrated intensity map. Right = Maximum intensity map. The source position is indicated by a black star. The beams for each map are shown in the bottom left corner, black for ^{13}CO and cyan for C^{18}O (Table 6.4). The contours used for C^{18}O here are the same as the contours used in Fig 6.7 (bottom row). The gas traced by these two molecules at the velocities quoted above show morphology of a similar size and direction.

6.3.4 C^{18}O (2–1) Kinematics

Fig. 6.6 (middle panel) shows the line profile of C^{18}O (2–1) at the innermost $0''.2$ from the source. This profile is quite similar in shape to the line profile of ^{13}CO (3–2). It

shows two components, with the blue side having a narrower and higher peak compared to the red side. As such a similar argument to this line profile as for ^{13}CO (3–2) is invoked. The C^{18}O (2–1) line profile may indicate there is gas associated to a disc in rotation. However, the double peaked shape of the C^{18}O (2–1) line profile may be caused by self absorption of warm photons closer to the central source being absorbed by cold gas in rotating envelope (Areal et al., 2018). As before, the exact shape of the line profile is probably due to mix of different emission features in the gas. The results of our line fitting are presented in Table 6.5.

Moving now to the velocity field maps for C^{18}O (2–1) in the two velocity channel regions (Fig. 6.7, right column). From $V_{lsr} = 5.7$ to 6.3 km s^{-1} . The gas seems to be slower closer to the source and faster as one moves out along the extension to the north-east. However, the slowest gas with respect to the systemic velocity is seen as a blob at $\geq 6.3 \text{ km s}^{-1}$ at around $0''.4$ in the north-west. Next, the velocity field map for the channels from $V_{lsr} = 6.9$ to 7.4 km s^{-1} are discussed. It is noted here that any changes in velocity seen in extended structures observed in the south-east and south-west are likely unreliable as the knot-like emission features along these directions appear smaller than the beam size. These knot-like emission features could be caused by the filtering of certain spatial scales by the ALMA interferometer in this high angular resolution array configuration. A flat extended structure around and to the west of the source of up to $\sim 0''.5$. Most of this structure is at a velocity of 7.15 km s^{-1} was seen but, its outer edge seems slightly faster. Although, the emission may be a superposition of different components.

Finally, two PV diagrams extracted from the C^{18}O (2–1) cube are shown. Two pseudo-slits were placed at PAs of 54° and 125° on the image cube. The velocity field map from $V_{lsr} = 6.9$ to 7.4 km s^{-1} is used for visualisation purposes (Fig 6.9, left panel). Two PV diagrams were extracted, one for each pseudo-slit. In Fig. 6.9 (top right panel) is the PV diagram extracted from the first pseudo-slit at 54° . Two key features are seen in this diagram. The first is some extended, non continuous emission that starts at $\sim -$

0".8 and is seen out to $\sim -2".4$. The velocity increases from $\sim 7.2 \text{ km s}^{-1}$ at $-0".8$ to $\sim 7.3 \text{ km s}^{-1}$ at $-2".4$. These positions indicate that this PV diagram is tracing emission from the south-western component observed in C^{18}O (2–1) from $V_{lsr} = 6.9$ to 7.4 km s^{-1} . The increase in velocity with distance from the source is referred to as a positive velocity gradient, which is indicative of accelerating material typical of the outflow launching process (Ferreira, 1997). Hence, this feature of the PV diagram shows that the south-western component is probably associated to out-flowing gas (Riaz et al., 2017; Whelan, 2014). The second feature is a slightly offset bump of $0".5$ at 7 km s^{-1} which might be part of gaseous emission from a circumstellar disc.

Fig. 6.9 (bottom right panel) presents the PV diagram extracted from the second pseudo-slit at 125° . Two features are also seen in this diagram. The first is a feature at a constant velocity of $V_{lsr} \sim 7 \text{ km s}^{-1}$ moving away from the source position from $0".4$ to $1".5$. This suggests the south-eastern component seen in C^{18}O (2–1) from $V_{lsr} = 6.9$ to 7.4 km s^{-1} may be part of an outflow cavity (Arce et al., 2013; Stahler, 1994). The second feature in this PV diagram is Keplerian-like (marked by the red arrows). This indicates there may be gas in rotation around the central source. However, it is noted that there is limited spatial and spectral resolution in these PV diagrams which inhibits a deeper analysis.

6.3.5 $\text{SO } 6_5\text{--}5_4$ Morphology

Figure 6.10 presents the integrated intensity (left panel) and maximum intensity (middle panel) maps for the $\text{SO } 6_5\text{--}5_4$ line. The emission traced by this line is much more compact compared to the gas traced by the CO isotopologues presented in the previous sections. The gas traced by $\text{SO } 6_5\text{--}5_4$ spans $< 0".4$ in any direction. Firstly, an emission of $\sim 0".2$ along the east-west direction and slightly south of the central source position is observed. A second component is seen to the north of the central source by about $0".25$. The maximum intensity map shows what appears to be a mix of these two structures. The peak contours in this map lie just to the west and north-east of the

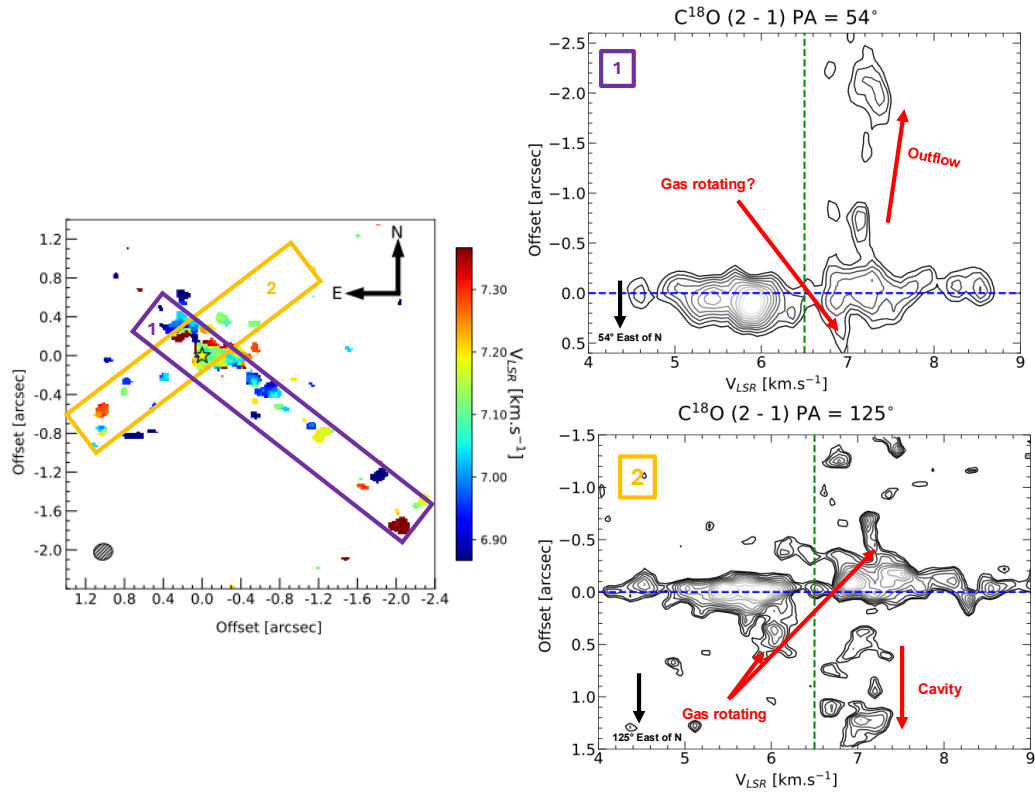


FIGURE 6.9: **Left:** The velocity map for $C^{18}O$ (2-1) from $V_{LSR} = 6.9$ to 7.4 km s⁻¹ (see Fig. 6.7) with the pseudo-slits overlaid for the corresponding PV diagrams. **Top right:** PV diagram extracted via slit 1 at PA = 54° east of north. Slit length = 3", slit width = 0".4 (20 pixels). The contours used were 4.7, 5.8, 7.0, 8.1, 9.3, 10.4, 11.5, 12.7, 13.8, 14.9 and 16.1 mJy/beam. **Bottom right:** PV diagram extracted via slit 2 at PA = 125° east of north. The slit was centred at the source position. Slit length= 3", slit width= 0".4 (20 pixels). The contours used were 4.3, 5.7, 7.0, 8.3, 9.7, 11.0, 12.3, 13.7, 13.8, 15.0, 16.4, 17.7 and 19.0 mJy/beam. A Gaussian smoothing filter has been applied to this PV diagram to increase the signal-to-noise. The green and blue dashed lines correspond to the systemic velocity (6.5 km s⁻¹) and zero offset from the source position respectively. The red arrows denote key features in each PV diagram.

central source. Above the north-east peak contour is a blob of emission about $0''.28$ north-east from the central source.

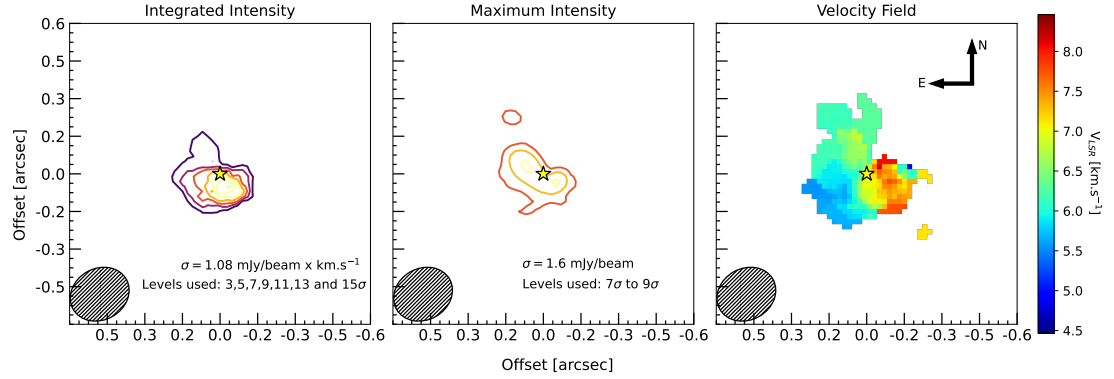


FIGURE 6.10: SO $6_5 - 5_4$ emission line moment maps. **Left:** Integrated intensity map from 4.5 to 8.5 km s $^{-1}$. Contours start at 3σ ($\sigma = 1.08$ mJy/beam $\times$ km s $^{-1}$) and increase in steps of 2σ . **Middle:** Maximum intensity map from 4.5 to 8.5 km s $^{-1}$. Contours start at 7σ ($\sigma = 1.60$ mJy/beam) and increase in steps of 1σ . **Right:** Velocity field map. The beam ($0''.24 \times 0''.20$ at PA = 119.4°) is represented as a black ellipse in the bottom left of each map. The source position is denoted by a yellow star.

6.3.6 SO $6_5 - 5_4$ Kinematics

In Fig. 6.6 the line profile for SO $6_5 - 5_4$ in the innermost $0''.2$ are shown. Like the line profiles of the CO isotopologues, this profile contains a blue and redshifted peak. However, unlike the line profiles of ^{13}CO (3–2) and C^{18}O (2–1) the intensity and FWHM of each component of the SO $6_5 - 5_4$ line profile are comparable. This may indicate that extended emission laying along the east-west direction could be related with a gaseous disc.

The right panel of Figure 6.10 presents the velocity field map of the SO $6_5 - 5_4$ emission. Like the integrated intensity map it shows the northern structure above the central source and the extended emission along the east-west direction. In the velocity field map, the emission in the east-west direction is blueshifted in the east and redshifted in the west. This is strong evidence that the gas detected in the east-west direction is associated with rotation. There is a hint of a change in velocity in the northern extended emission. The top of the emission at $\sim 0''.3$ has a velocity of $V_{lsr} \sim 6.2$ km s $^{-1}$.

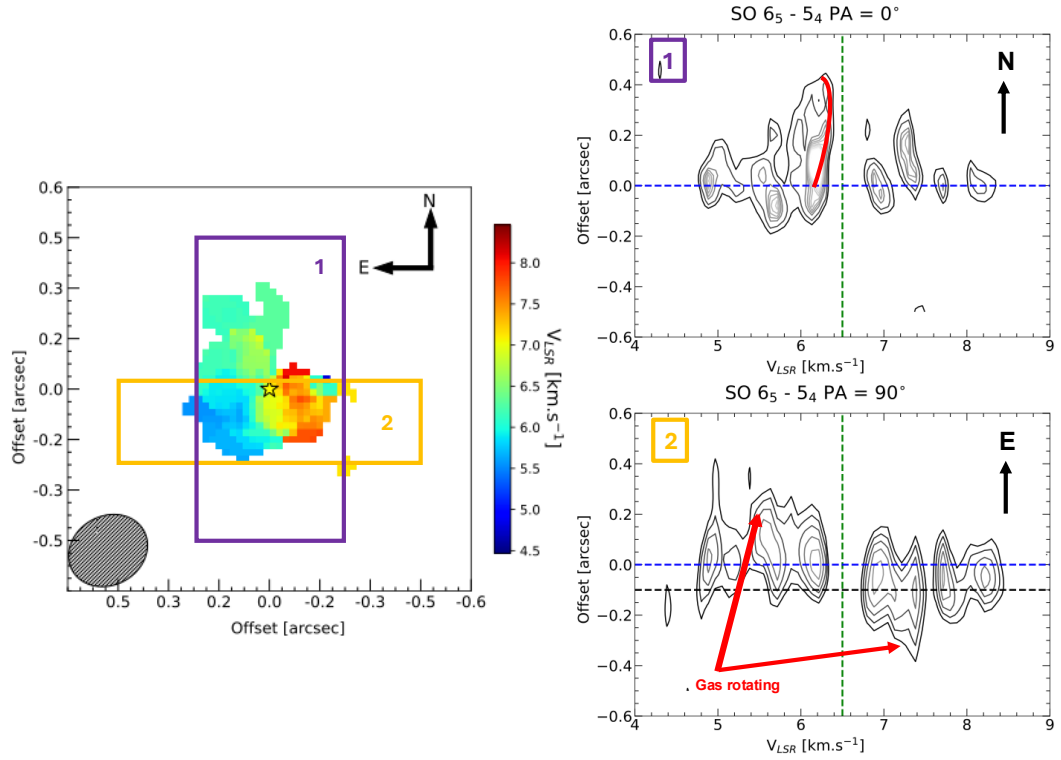


FIGURE 6.11: **Left:** The velocity map for the SO $6_5 - 5_4$ emission line (see Fig 6.10 right panel). **Top right:** PV diagram extracted via slit one at $PA = 0^\circ$ east of north. The slit was centred at the source position. Slit length= 1", slit width= 0".42 (21 pixels). The contours used were 6.5, 7.9, 9.4, 10.0, 10.4, 10.8, 11.2, 11.5, 11.8, 12.0, 12.3 and 12.5 mJy/beam. A red arc has been drawn by eye to aid in visualisation of an observed velocity gradient in this PV diagram. This gradient may indicate a possible path for the streamer in position-velocity space. **Bottom right:** PV diagram extracted via slit two at $PA = 90^\circ$ east of north. The slit was centred 0".1 below the source position to encapsulate as much of the emission along this PA as possible. Slit length= 1", slit width= 0".24 (12 pixels). The contours used were 5.1, 6.3, 7.5, 8.7, 9.9, 11.0, 12.2 and 13.4 mJy/beam. The black dashed line corresponds to the centre position of the slit. The green and blue dashed lines correspond to the systemic velocity (6.5 km s $^{-1}$) and zero offset from the source position respectively.

At the bottom close to the position of the central source the velocity changes to $V_{lsr} \sim 6 \text{ km s}^{-1}$ (i.e. the gas seems to become more blueshifted with respect to the systemic velocity as one moves closer the central source).

Fig. 6.11 presents the PV diagrams extracted from the SO 6_5-5_4 emission cube. The PV diagram extracted via the first pseudo-slit is shown in the top right panel of Fig. 6.11. The slit was placed along the north-south direction with the aim of capturing the emission from the northern extended structure. The outer contour in this PV diagram

which extends to an offset of $\sim 0''.425$ is at $V_{lsr} \sim 6.2 \text{ km s}^{-1}$. As one moves closer to the central source the velocity becomes more blueshifted with respect to the systemic velocity. The velocity decreases to $V_{lsr} \sim 6 \text{ km s}^{-1}$ at the position of the source. The change in velocity of the northern extended emission closer to the source suggests that the gas is speeding up towards the source. Also, considering that this northern emission is only extended from the central source by $\sim 0''.3$, hints that this material may be gravitationally bound. These two pieces of information suggest the possible presence of part of an accretion streamer which may be impacting a circumstellar disc.

The PV diagram extracted via the second pseudo-slit is shown in the bottom right panel of Fig. 6.11. The slit was placed along the east-west direction. This PV diagram shows that the blueshifted side at 5.4 km s^{-1} has a positive offset from the central source of $\sim 0''.3$ and the redshifted side at of the disc at 7.4 km s^{-1} has a negative offset of $\sim -0''.4$. This gradient in velocity close to the central source is consistent with gas rotating anti-clockwise.

6.4 Discussion

From this analysis and the literature, IRAM04191 has observational properties that are consistent with a star-like formation scenario. In the literature, at large spatial scales, there is evidence of an outflow (André, Motte, and Bacmann, 1999), outflow cavity (Lee, Ho, and White, 2005; Lee et al., 2002) and a large rotating natal envelope (Belloche et al., 2002). At small spatial scales, our analysis provides evidence of both mass-loss and mass accretion processes in the form of outflows and/or outflow cavities (e.g. Fig. 6.5 and Fig. 6.8) and a potential accretion streamer observed in $\text{SO } 6_5-5_4$ (e.g. Fig. 6.11 top right panel) respectively. The many extended structures in the $^{13}\text{CO } (3-2)$ and $\text{C}^{18}\text{O } (2-1)$ seem to be tracing shocked gas and a potential second outflow in the case of the blueshifted $^{13}\text{CO } (3-2)$, and a cavity in the case of the redshifted $^{13}\text{CO } (3-2)$ and $\text{C}^{18}\text{O } (2-1)$ emission.

As mentioned previously, small scale streamers have been observed in SO 6₅–5₄ before with a few examples observed in Class 0/I YSOs (Flores et al., 2023; Kido et al., 2023) where these streamers are modelled as either in-falling rotating structures or as material impacting a flattened disc like structure. The observations presented here appear to suggest this second scenario for the potential streamer given its close proximity to the disc like structure seen in SO 6₅–5₄. However, this proximity between the two structures may be a projection effect. Accretion streamers have been of particular interest in recent years as they may (at least partly) explain the phenomena of episodic accretion observed in YSOs (e.g. EX Lupi). They have been observed around several Class 0,I and II YSOs (Hanawa et al., 2024; Hsieh et al., 2023; Pineda et al., 2020; Valdivia-Mena et al., 2024; Yen et al., 2019), and have been detected previously, with a handful of carbon based molecular tracers (e.g. HCO⁺, HC₃N, DCN, CS, and C¹⁸O) and in 1.3 millimetre continuum. The literature has also noted that SO may trace the inner region of an accretion streamer close to the Class II source, HL Tau (see Garufi et al. (2022), their Fig. 3, panel a).

One leading model of BD formation purport that BDs should form as a scaled down version of low-mass stars (see Palau et al., 2024, and references therein). As such, one would expect to observe shorter, slower and less powerful outflows, lower mass accretion rates and smaller streamers of in-falling material in very young BDs. A comparison between what has been observed for both low mass protostars and young BDs is now discussed.

As described in Chapter 1, the formation process of low mass stars in the Class 0/I stage is characterised by both mass accretion and mass-loss processes. Molecular outflows of length 0.1 to 5 pc with velocities in the order of $\sim 10 \text{ km s}^{-1}$ have been observed in a variety of tracers. They are interpreted as material swept up by the faster, inner molecular jet (Bally, 2016; Frank et al., 2014). Episodic accretion bursts have been detected in the Class 0/I stage for stellar masses as low as $0.1 M_{\odot}$ (Stock et al., 2020). These bursts are generally associated to interactions with disc material. Streamers of

in-falling material interacting with the circumstellar disc may play a part in explaining the origin episodic accretion. They have been observed for a handful YSOs at a variety of spatial scales (both in and outside the natal envelope) at the Class 0/I stage (Hales et al., 2024; Pineda et al., 2020; Valdivia-Mena et al., 2024).

In comparison, for BDs, Whelan et al. (2005) and Riaz et al. (2017) report the first atomic outflows launched by a BD and proto-BD respectively. The projected lengths and speeds of these reported outflows are $0''.08$ to $0''.1$ (10 to 12.5 AU at the assumed distance of 125 pc to ρ -Ophiuchi), at a blueshifted velocity of approximately -40 km s^{-1} and 0.26 pc ($\sim 5.36 \times 10^4 \text{ AU}$ at an assumed distance of $387 \pm 1.3 \text{ pc}$ to the driving source M1701117) at a blueshifted velocity of $\sim -45 \text{ km s}^{-1}$ respectively. In terms of accretion processes, with particular emphasis on streamers, Riaz, Stamatellos, and MacHida (2024) report the only known detection of an accretion streamer in the vicinity of a proto-BD candidate to date. The reported projected length of this streamer is 2000 to 3000 AU. This is smaller in scale to the streamers seen in the vicinity of low mass protostars (e.g. Pineda et al., 2020).

The potential streamer of gas detected in SO 6_5-5_4 close to IRAM04191 is of projected length of $0''.3$ (42 AU at the assumed distance to IRAM04191). Also, considering the length scale of the known CO outflow detected by André, Motte, and Bacmann (1999), which seems more typical of outflows launched by low-mass protostars, and the small projected size of the potential streamer, IRAM04191 shows observational characteristics of both low-mass YSOs and BDs. Lastly, the flattened structure in SO 6_5-5_4 along the east-west direction of IRAM04191 has similar velocity structure and morphology to the Keplerian disc detected around this source in C^{18}O (2–1) (Maury, 2020).

The results presented in this chapter have mixed indications. There is evidence that may support the presence of an accretion streamer and hence, possible episodic accretion. This could support the idea that IRAM04191 could end up as a stellar mass object by the end of the Class 0/I stage, if it accretes enough mass to reach the stellar regime.

However, it is currently unknown how much mass can be gathered through these accretion streamers. Streamers may be an important source of material for young forming stars and BDs as the bulk of a low mass YSOs final mass is accreted in the Class 0/I stage (Fiorellino et al., 2023). Alongside this, evidence of large and small scale mass-loss processes (outflows and outflow cavities) add further complication to the mass assembly picture for IRAM04191. As such, accretion streamers and their relationship to outflows will be an important avenue of further study. Given the above, conclusions on the fate of the IRAM04191 system cannot be made as more observations are needed for three reasons.

1. Further observations are required at high spatial and spectral resolution to map gas at the spatial scales in-between the results presented here, and the large scale structures reported in the literature (André, Motte, and Bacmann, 1999; Lee, Ho, and White, 2005; Lee et al., 2002). These proposed observations will be critical to assess the link between large and small scale mass-loss and mass accretion processes.
2. High spatial resolution dust continuum observations are needed in the cm and mm/sub-mm to obtain a reliable mass estimate for the central source. This will confirm if IRAM04191 is currently a sub-stellar or stellar mass object and may also shed light on its possible binarity (Lee, Ho, and White, 2005). These observations and analysis have been provided by Huélamo et al. (2026). These authors use a combination of archival ALMA sub-mm observations and cm observations from the Very Large Array to confirm the presence of a gaseous disc traced in C¹⁸O (2–1) gas. Additionally, continuum observations suggest the presence of two objects in the IRAM04191 system. A Keplerian curve fit to the PV diagram extracted along the major axis of the C¹⁸O (2–1) disc puts a dynamical mass constraint of $50 \pm 50 M_{\text{Jupiter}}$ on the IRAM04191 system. Therefore, these combined observations suggest that IRAM04191 is a proto-BD binary candidate system.
3. A monitoring campaign with observations taken at a high cadence in time are

needed to confirm if this source goes into outburst because of episodic accretion events. Other VeLLOs, for example IRAS 15398–3359, have also been suggested to be out-bursting (Jørgensen et al., 2013). If these outbursts are confirmed it will provide strong evidence that IRAM04191 may evolve into a low-mass star by the end of the Class 0 stage. These observations may also tell us if these outbursts are EXor like or FUor like (Fischer et al., 2023).

6.5 Conclusions and Summary

A detailed morphological and kinematic study of high density gas in the immediate proximity of IRAM04191 as traced by three emission lines ^{13}CO (3–2), C^{18}O (2–1) and SO 6₅–5₄ has been carried out. The summary and conclusions of this study are as follows:

1. Emission from ^{13}CO (3–2) is complex. The blueshifted emission traces gas on larger spatial scales compared to C^{18}O (2–1) or SO 6₅–5₄. The ring-like structure in the blueshifted gas overlaps with the arc of gas traced by the 4.5 μm passband to the north-west of the Spitzer/IRAC imaging. Hence, this ring-like structure may be tracing previously reported shocked gas. This is interesting as it may be an indication of another outflow in this direction, suggesting a possible binary nature for the source as pointed out in previous works.
2. The redshifted ^{13}CO (3–2) gas is also complex. The spurious emission to the north-west may also be tracing shocked gas. Whereas, the structure to the south-east may be tracing the base of an outflow cavity.
3. A velocity dispersion gradient is detected in the innermost 2" of the blueshifted ^{13}CO (3–2) gas. This may indicate an increase in turbulence in the gas close to the source. This can be caused by launching of a jet/outflow or due to the impact of material in-falling onto the circumstellar disc.
4. The emission from the C^{18}O (2–1) line is also complex. It shows several emission components at different PAs. One of these components traced in the velocity

channels from $V_{lsr} = 6.9$ to 7.4 km s^{-1} , has morphology consistent with an even denser part of an outflow cavity (Fig. 6.8). The PV diagram of this emission component (Fig 6.9, right panels) shows kinematic signatures associated with out-flowing gas. The other emission components detected in the other velocity channels seem to have a mixed origin. Some may be associated to the outflow found by Podio et al. (2021) and some may be linked to disc emission or may be part of a rotating envelope.

5. A gas disc seems to be present around IRAM04191, based on the double peaked profiles of ^{13}CO (3–2), C^{18}O (2–1) and SO 6₅–5₄. The PV diagram in SO 6₅–5₄ shows a velocity gradient in the east-west direction consistent with gas rotating anti-clockwise. This rotation direction is opposite to that of the rotating envelope and hence, may strengthen the argument that this structure is a forming gas disc (Maury, 2020). The PV diagram in C^{18}O (2–1) at a PA of 125° also shows a signature that may hint towards rotating gas. It is noted that for both PV diagrams that a Keplerian profile cannot be determined with the current angular resolution of the presented data.
6. A potential streamer of in-falling material towards IRAM04191 is also detected to the north of the central source in SO 6₅–5₄ (Fig. 6.10 right panel). This line is a known streamer tracer (Flores et al., 2023; Garufi et al., 2022; Kido et al., 2023) from Class 0 up to Class II low mass YSOs. The PV diagram extracted along the north-south direction (Fig. 6.11 top right panel) shows a kinematic signature consistent with in-falling gas. This is interesting given that it is perpendicular to the structure observed by Maury (2020) and on a comparable spatial scale. This could be a hint that this potential streamer may be interacting with the flattened disc like structure that was previously observed.

In conclusion, the results presented and discussed above clearly show that IRAM04191 has a complex formation history. There is evidence suggesting that both mass accretion and mass-loss processes are occurring simultaneously in this system. From this, the results presented in this study suggest that the formation pathway for IRAM04191 seems to be star-like rather than planet-like but, conclusion on the fate of this object

post the Class 0 stage cannot be made. In this context, to understand if the majority of BDs are formed as stars, more substellar candidates in very early formation stages need to be studied in detail. VeLLOs such as IRAM04191 may prove to be promising sub-stellar candidates and hence warrant further study. Future JWST observations will be critical in finding and characterising these VeLLO objects to determine if they are sub-stellar.

Chapter 7

Overall conclusions and future work

Outflows are ubiquitous at all mass scales and likely play a pivotal role in extracting angular momentum from the stellar and sub-stellar systems from which they are launched. In doing so, it is thought that outflows are the mechanism that enables accretion onto the central source. Alongside this, they provide an essential source of feedback into the natal environment in which these objects are born and may even help "kickstart" a new generation of star formation. Additionally, these flows impact the physical and chemical properties of the discs of dust and gas surrounding young stars and BDs therefore, effecting the initial conditions of planet formation. All these questions regarding outflows are critical to understanding our origin and place in the Universe and hence, make the study outflows an exciting topic in modern astrophysics.

This thesis has explored the role that outflows play during the formation process of young stars and BDs. Utilising the forefront observational techniques of optical spectro-astrometry and radio interferometry, the presented studies have explored the role outflows play during the assembly of YSOs and BDs. This concluding chapter summarises the findings and contributions of this thesis, suggest improvements and expansions to the current analysis, and discuss avenues of future research in the fields of star and BD formation.

7.1 Summary of findings and contributions

7.1.1 Understanding of the optical FEL low velocity component.

Chapter 4 presented a spectro-astrometric analysis of ten CTTSs in the [O I] $\lambda 6300$ and [S II] $\lambda 6731$ FELs in an attempt to better understand the origin of the optical FEL LVC and its role in the angular momentum transport in young star-disc systems. Additionally, the analysis presented in this chapter aimed to provide an observational test to one of the key assumptions of disc wind driven accretion models; namely, that the mass loss rate of the low velocity wind should be comparable to the stellar mass accretion rate. In performing this analysis and attempting to answer these questions, Chapter 4 is a direct extension to the pilot studies of Whelan et al. (2021) and Birney et al. (2024b). These studies combined show that in the case of one CTTS, RU Lupi, the LVC-NC has its origin in a slow an MHD disc wind launched from within the gravitational well of the star (Whelan et al., 2021) and that this wind is inefficient at removing excess angular momentum (Birney et al., 2024b), in contradiction to the assumption made by models.

The origin of the LVC emission in these sources could have various origin. Either arising from the circumstellar disc itself, the base of a disc wind or (as was the case of RU Lupi), the [O I] $\lambda 6300$ LVC may be a transitional layer of the outflow system between the denser atomic regions and a more diffuse molecular part of the outflow (Birney et al., 2024b; Nisini et al., 2024). The results presented in Chapter 4 appear to indicate that the LVC emission in this larger sample of CTTSs is very spatially compact. This evidence hints that the LVC emission in these CTTSs likely arises from a compact region of the outflow system close to the stars (i.e. potentially the wind base) or from trapped gas within (or just above) the disc itself. However, the mass outflow rates and outflow efficiency estimates for this sample of ten CTTSs studied in Chapter 4 vary widely between different stars due to poor constraints on the temperature of the material being traced by the optical LVC (i.e. the temperature is anywhere between 5,000 and 10,000 K). This limits this studies ability in testing the assumption of wind driven

accretion models. Given these findings, it leads to two primary conclusions. Firstly, the LVC-NC of RU Lupi may be unique in tracing a disc wind and therefore the overall origin of the optical LVC-NC remains unknown. Secondly, the [O I] $\lambda 6300$ FEL may not ideal constraining disc wind mass fluxes. This is in agreement with the findings of Nisini et al. (2024) and leads to the assertion that new disc wind tracers need to be studied in order to constrain the temperature and mass flux of the outflowing material. Overall, the work presented in Chapter 4 provides interesting finding in regard to the optical LVC.

Chapter 5 also attempts to understand the origin of the optical LVC but via a different approach. The aim of this chapter was to examine the spatial and spectral properties of the optical FEL LVC of the CTTS DG Tau over several epochs of high spectral resolution data spanning ~ 20 years, given the presence of a known variable jet; traced by the HVC. The findings presented in Chapter 5 show that the optical FEL profiles of DG Tau are complex and vary in shape over time. This is due to varying contributions to the line flux from different line sub-components between the three epochs of data. Despite this, the spectro-astrometric analysis of the LVC of DG Tau remains remarkable consistent. A negative velocity gradient was observed in all three of the studied FELs in all epochs for the LVC-M sub-component. This is a clear indication of this LVC sub-components' origin in a disc wind (similar to RU Lup). Additional findings from this chapter are in agreement with the existing literature. For example, there is clear evidence that the high velocity jet of DG Tau has been slowing over the ~ 20 years of time that these observations are spread. This is in agreement with the works of Takami et al. (2023) and Pyo et al. (2024). Overall the work presented in this chapter provides some interesting constrains on the origin of the LVC of DG Tau (i.e. that it originates in a disc wind) but, in addition provides some interesting constraints for disc wind models in general; with results indicating that the LVC may not change in response to a variable jet. However, the possibility of a time delay between a variable jet and corresponding changes in the LVC needs further exploration.

7.1.2 Sub-stellar outflows and accretion streamers

Chapter 6 presented ALMA observations of high density gas surrounding the VeLLO system IRAM 04191+1522. The aim of this work was to uncover the formation "pathway" of this source and to learn more about the role outflows play in the early phases of BD formation. This study revealed a highly complex environment shaped by both accretion and outflow processes. The ^{13}CO (3–2) emission traces large scale (~ 1000 AU), shock related structures possibly linked to multiple outflows. The detection of a second outflow at a different PA to the primary molecular flow observed in CO adds further to the potential binary nature for this system. A binary nature for this source has recently been suggested by new continuum observations presented by Huélamo et al. (2026). If a second outflow is present, this places some interesting constraints on current models of BD formation and firmly places IRAM04191 along the "star-like" formation pathway. Given this, it appears that outflows play a comparable role at sub-stellar masses in regulating the formation process of these young forming objects.

In addition to the detection a potential second outflow, this chapter also provides evidence which hints at presence of a collapsing gas disc surrounding IRAM04191. This is supported by both double peaked line profiles in all three molecular lines and the SO 6_5-5_4 velocity field map which indicates that this structure is rotating opposite to the rotation direction of the envelope. Although, it is noted that current spectral and spatial resolution limitations prevent confirmation of a Keplerian motion pattern. However, this Keplerian rotation has been confirmed in C^{18}O by Huélamo et al. (2026). The SO 6_5-5_4 emission also reveals a possible infalling streamer interacting with the gas disc in close proximity (few 10s of AU) to IRAM 04191+1522. Overall, this system is highly complex showing signatures of ongoing mass accretion and mass loss in addition to likely binarity. This combination of observed features is uncommon in BD systems. As such IRAM04191 will be an exciting laboratory for future studies of BD formation.

Overall, the work presented in these three results chapters show that outflows play a

critical role in the star and BD formation processes. Regulating angular momentum transport and enabling mass accretion. These studies have placed valuable constraints on the current models while simultaneously opening up new questions in the field. The next section will discuss these questions and avenues for future research.

7.2 Future work

7.2.1 Extension to the spectro-astrometric survey

In the short term, an extension the analysis presented in Chapter 4 is planned. Additional data has been obtained to expand the original ten targets presented in Chapter 4 to a further set of six CTTs. UVES observations were taken in period 112 (PI: N. Otten, Programme ID: 112.25P0.001) and will be analysed in a future article. The end goal is to eventually obtain a statistically significant sample (i.e. ≥ 30) of stars with varying properties over different star forming regions to confirm what exactly the LVC-NC in optical FELs traces in the majority of CTTs.

7.2.2 The future of spectro-astrometry

The future of SA also looks bright. In the upcoming era of the European Southern Observatorys' (ESO's), Extremely Large Telescope (ELT) there are predictions that it could be feasible for SA to detect solar system like exomoons at moderate temperatures (150 - 300K, van Woerkom and Kleisioti, 2024). Alongside this, recent numerical simulations suggest that future waveguide technology called photonic lanterns (PLs) could be used in future fibre-fed spectrographs to enable the recovery of spectro-astrometric offsets without the need for multiple slit PAs (Fig. 2.1) as required by traditional echelle or long-slit spectrographs (Kim et al., 2024). Additionally, advanced space telescopes such as JWST could also make use of SA through its near infrared spectrograph (NIRSpec).

7.2.3 New disc wind tracers

As noted by the works of Nisini et al. (2024) and Birney et al. (2024b); the optical [O I] $\lambda 6300$ line is likely not the best line to test angular momentum removal given uncertainty in the temperature of the material traced by this line and the variation in mass flux observed in the sample of CTTs presented in Chapter 4. The work of this thesis reaffirms the conclusion of Nisini et al. (2024) regarding the [O I] $\lambda 6300$ line. The work presented here suggests that cooler molecular lines such as H₂ or CO may be more suitable for tracing low velocity disc winds. Future studies could involve spectro-astrometric and spectro-imaging observations of infrared H₂ lines using VLT/CRIRES (CRyogenic high-resolution InfraRed Echelle Spectrograph) and VLT/ERIS (Enhanced Resolution Imaging Spectrograph), respectively, alongside ALMA observations of the CO gas. These observations could be used to constrain the properties (e.g. mass flux, velocity, length, collimation etc.) of the outflow structures that these lines may trace. JWST may also offer interesting insights into outflows driven by young stars via observations with NIRSpec as shown by Pascucci et al. (2024) and in mid infrared emission using the Mid InfraRed Instrument (MIRI).

7.2.4 Multi epoch high cadence data

It is now clear that multi epoch high time cadence observation are required to understand low velocity winds and their relationship to the high velocity jets. The results and conclusions in Chapter 5 emphasise this. In conjunction with this, monitoring campaigns to study the spectral and spatial variability of both the accretion and ejection processes in young low-mass stars are needed to answer questions about how these processes affect the star and planet formation process. This is important given the many different kinds of variable phenomena are observed in low mass stars (Fischer et al., 2023).

Dedicated multi epoch imaging, spectral and interferometric observations and with

various instruments (e.g. VLT/X-Shooter, VLT/MUSE and the telescopes and instruments mentioned above) are now required to understand the impact that variable accretion (e.g. accretion bursts, flares etc.) and ejection has on the formation processes of stars, BDs and planets over time. More specifically, some questions that arise from variability come from how this affects the mass assembly process in stars and BDs, the evolution and dispersal of the disc, and how it changes the physical and chemical conditions of the disc during the formation of planets. Variability may also play a role in how stellar and sub-stellar outflows affect the surrounding environment.

7.2.5 Detailed studies of accretion streamers

Accretion streamers which have been recently discovered around low mass YSOs and BDs (Kim et al., 2019; Pineda et al., 2020; Riaz, Stamatellos, and MacHida, 2024; Riaz et al., 2017) now puts the classic picture of isolated star formation into question. These structures directly impact the discs around YSOs and BDs and may supply fresh material available to be accreted onto the central source. Characterisation of these structures is needed to fully understand the impact they have on the star formation process. However, the limited number of observations of these structures in just a handful of molecular emission lines makes it difficult to ascertain how much mass is delivered to the disc and how the impact of these structures affects the star-disc system. Future observation with ALMA should now focus on finding as many appropriate streamer tracing lines as possible with the aim of detecting more of these structures around YSOs and BDs. Alongside these ALMA observations, complementary observations in the mid-infrared with MIRI on JWST could help constrain the physical structure of streamers (i.e. what molecule traces which part of the accretion streamer), help to assess the impact they have on discs, constrain the mass of the streamers, and the mass flux delivered by them. Additionally, these observations could shed light on the role streamers have in regard to variable accretion and accretion bursts in YSOs and BDs.

7.2.6 The distinction between disc winds and accretion streamers

One final question this thesis raises is in regard to the distinction between disc winds and small scale (or structures within) accretion streamers, like the potential streamer observed in SO emission in Chapter 6. Disc winds show a negative velocity gradient where the wind streamlines decrease in velocity with height above the disc mid plane. In the case of accretion streamers, the gravitationally bound material accelerates towards the star-disc system is such that the material in the part of the streamer closest to the disc is faster than the part of the streamer further above the disc. Given this, streamers may also show a similar velocity gradient to that typical of disc winds. Usually, the distinguishing factor between accretion streamers and disc winds is morphology. Disc winds appear to be conical in their morphology, whereas large scale streamers tend to have complex morphological structure with curves (e.g. Pineda et al., 2020). At small scales close to the disc, this distinction in morphology between accretion streamers and disc winds is unclear. Future facilities such as ESOs' ELT, the wideband sensitivity upgrade for ALMA and the next generation Very Large Array (ngVLA) observing in the optical/near infrared, the mm-sub-mm, and centimetre wavebands, respectively, may offer instruments capable of performing the observations needed to disentangle these processes.

7.3 Final remarks

The work presented in this thesis provides some useful insights into the formation processes of YSOs and BDs. In particular, this work gives some interesting constraints in regard to the outflow launching models and role of outflows in both the sub-stellar and stellar mass regimes. There are still many open questions regarding outflows in these systems. Some of the observations proposed above could shed light on these unanswered questions. In addition, the role of accretion streamers and variability in both the accretion and outflow processes in these young forming systems is opening a new frontier in the fields of star, BD and planet formation. Tackling these questions will require new technology and new approaches in observational strategy and data

analysis. Ultimately, answering these fundamental questions about the mechanism(s) that create these systems will bring us closer to understanding our own origins and place within the Universe.

Appendix A

Observational log for the period 89 and 108 runs

TABLE A.1: Observing log for the period 89 and 108 runs.

Target	RA [J2000]	Dec [J2000]	Date [dd-mm-yyyy]	Slit PA [°]	Exposure Times [ncycles × s]
TW Hya	11:01:51.9	-34:42:17.03	02-05-2012	62	5 × 300
			01/02-05-2012	152	10 × 300
			02-05-2012	242	5 × 300
			01/02-05-2012	332	10 × 300
T Cha	11:57:13.6	-79:21:31.5	02-05-2012	67	5 × 300
			02-05-2012	157	5 × 300
			02-05-2012	247	5 × 300
			02-05-2012	337	5 × 300
MP Mus	13:22:07.5	-69:38:12.2	29-06-2012	169	3 × 300
			29-06-2012	349	3 × 300
MP Mus (0".3 slit)	13:22:07.5	-69:38:12.2	02-05-2012	79	3 × 300
			02-05-2012	259	3 × 300
LKHa 330	03:45:48.28	+32:24:11.85	02-11-2021	75	1 × 300
			02-11-2021	165	1 × 300
			15-11-2021	75	3 × 900
			15-11-2021	165	3 × 900
			15-11-2021	255	2 × 1000
			16-11-2021	345	2 × 1000
GQ Lup	15:49:12.11	-35:39:05.058	19-02-2022	76	1 × 900
			14-03-2022	256	1 × 1200
			01-04-2022	76	2 × 1200

Continued on next page

Target	RA [J2000]	Dec [J2000]	Date [dd-mm-yyyy]	Slit PA [°]	Exposure Times [ncycles × s]
DG Tau	04:27:04.69	+26:06:16.06	02-04-2022	166	2 × 1200
			04-04-2022	256	2 × 1200
			14-11-2021	46	1 × 300
			14-11-2021	136	1 × 300
			16-11-2021	46	3 × 800
CW Tau	04:14:17.00	+28:10:57.766	13-02-2022	136	3 × 800
			02-11-2021	54	1 × 300
			14-11-2021	54	1 × 300
			02-11-2021	144	1 × 300
			14-11-2021	144	1 × 300
GK Tau	04:33:34.56	+24:21:05.86	10-11-2021	234	1 × 1000
			02-02-2022	234	2 × 1000
			16-02-2022	324	2 × 1000
			04-11-2021	300	2 × 1200
			06-11-2021	210	1 × 1000
			15-11-2021	30	2 × 1050
			15-11-2021	300	1 × 1000
			28-11-2021	30	1 × 1000
			02-02-2022	120	1 × 1000
			16-02-2022	120	2 × 1050
DK Tau	04:30:44.24	+26:01:24.65	27-01-2022	4	1 × 1600
			05-02-2022	4	1 × 1600
			14-02-2022	4	2 × 1300
			17-02-2022	184	2 × 1450
			18-02-2022	274	2 × 1450
EX Lup	16:03:05.49	-40:18:25.43	03-10-2021	243	3 × 600
			30-01-2022	243	2 × 600
			13-02-2022	243	2 × 600
			13-02-2022	63	4 × 600
			13-02-2022	153	4 × 600
			04-03-2022	243	3 × 600
			14-03-2022	333	2 × 600
			29-03-2022	333	3 × 600

Appendix B

Channel Maps for the molecular emission line cubes

This appendix presents the channel maps for the ^{13}CO 3–2, C^{18}O 2–1 and SO 6₅–5₄ molecular emission lines that were analysed and discussed in Chapter 6.

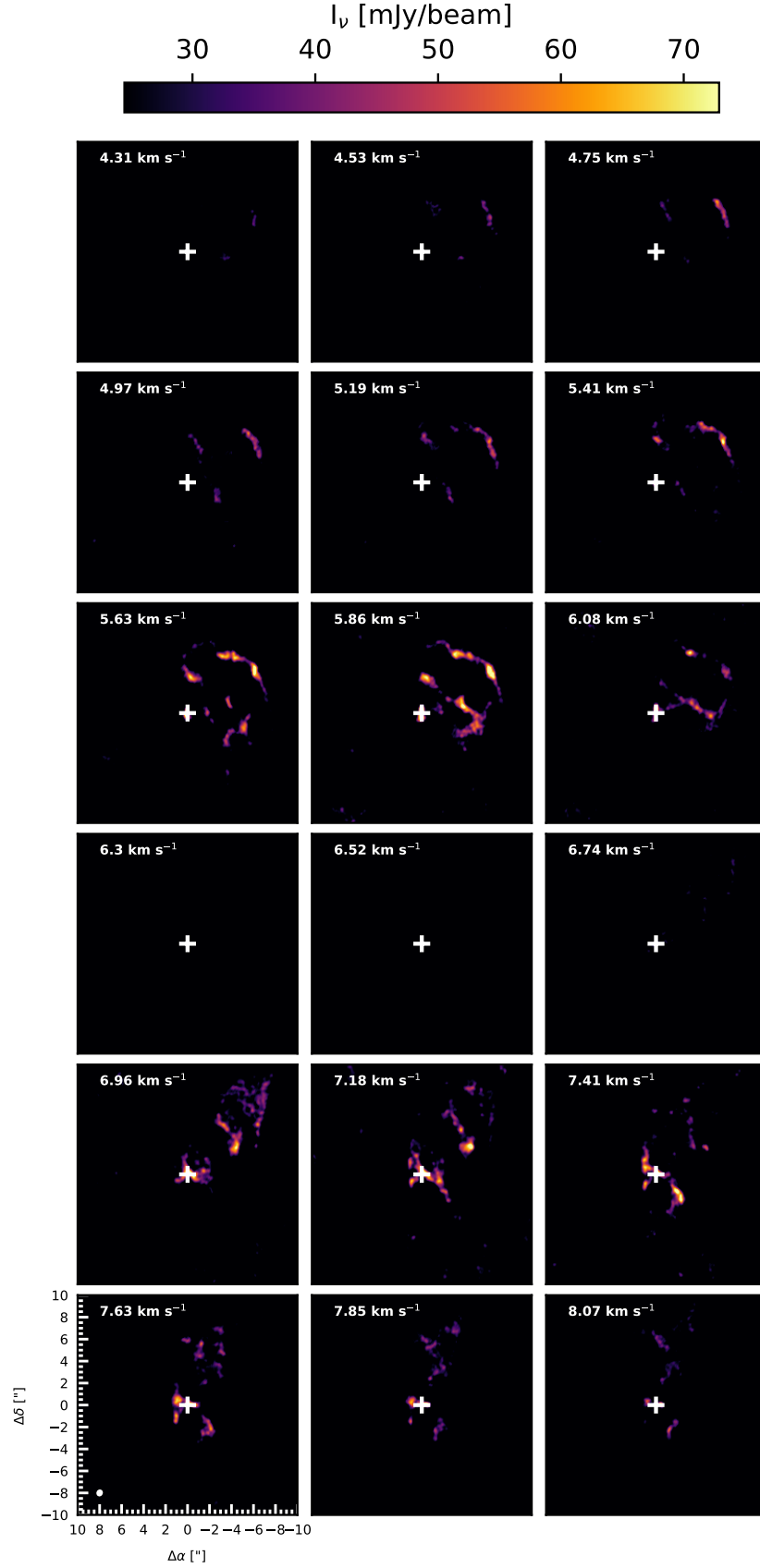


FIGURE B.1: Channel maps for ^{13}CO 3–2 emission in IRAM04191+1522 from $V_{\text{LSR}} = 4.31 \text{ km s}^{-1}$ to 8.07 km s^{-1} . The emission shown is clipped at $3\sigma_{\text{rms}}$ and above ($\sigma_{\text{rms}} = 9.484 \text{ mJy/beam}$). The beam ($0''.39 \times 0''.33$ at 155.6°) is shown in the bottom left of the lower left channel map. The systemic velocity is 6.5 km s^{-1} . The source position is indicated by the white cross.

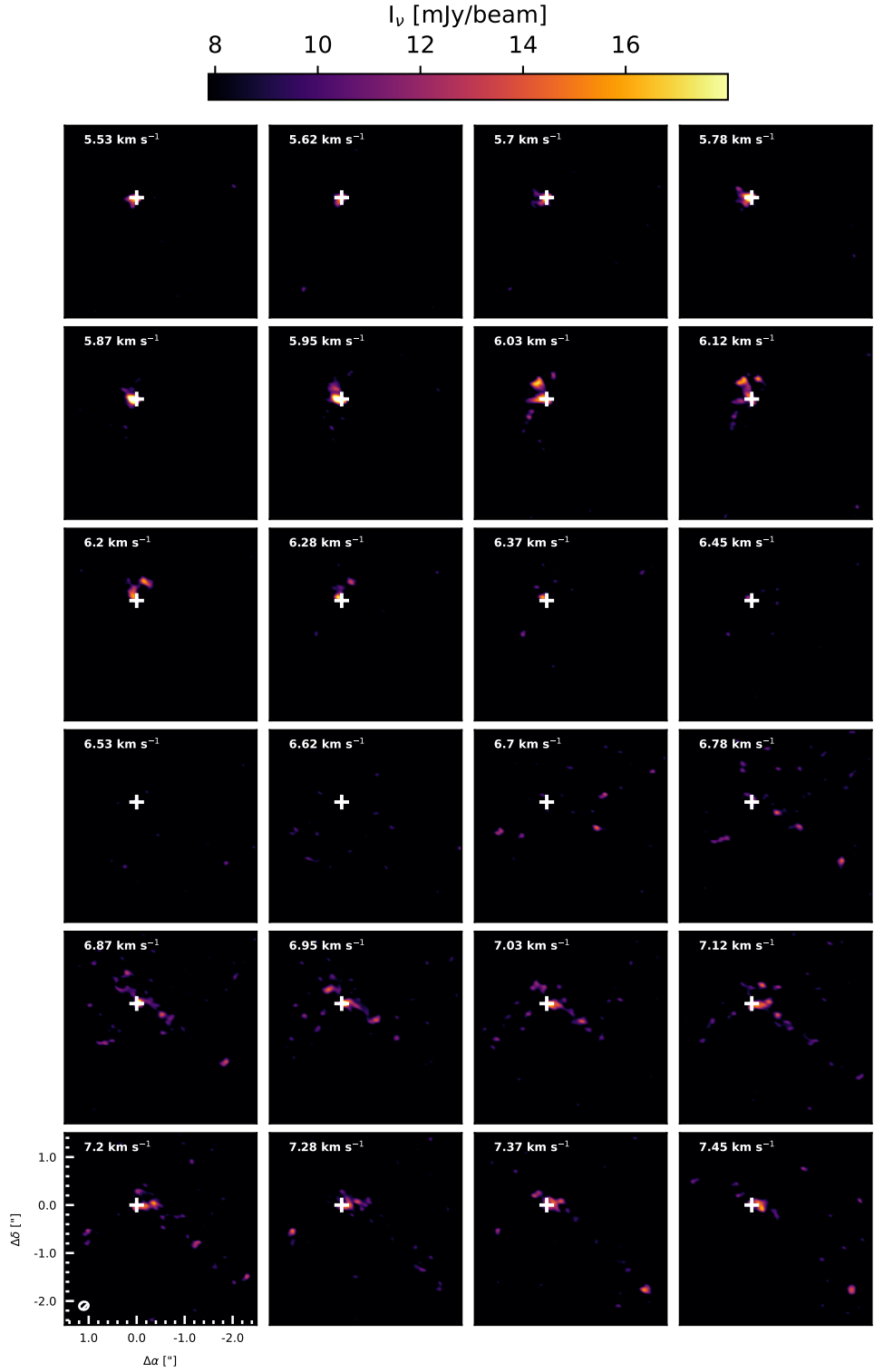


FIGURE B.2: Channel maps for C^{18}O 2–1 emission in IRAM04191+1522 from $V_{\text{LSR}} = 5.53 \text{ km s}^{-1}$ to 7.45 km s^{-1} . The emission shown is clipped at $3\sigma_{\text{rms}}$ and above ($\sigma_{\text{rms}} = 2.623 \text{ mJy/beam}$). The beam ($0''.19 \times 0''.17$ at 102.1°) is shown in the bottom left of the lower left channel map. The systemic velocity is 6.5 km s^{-1} . The source position is indicated by the white cross.

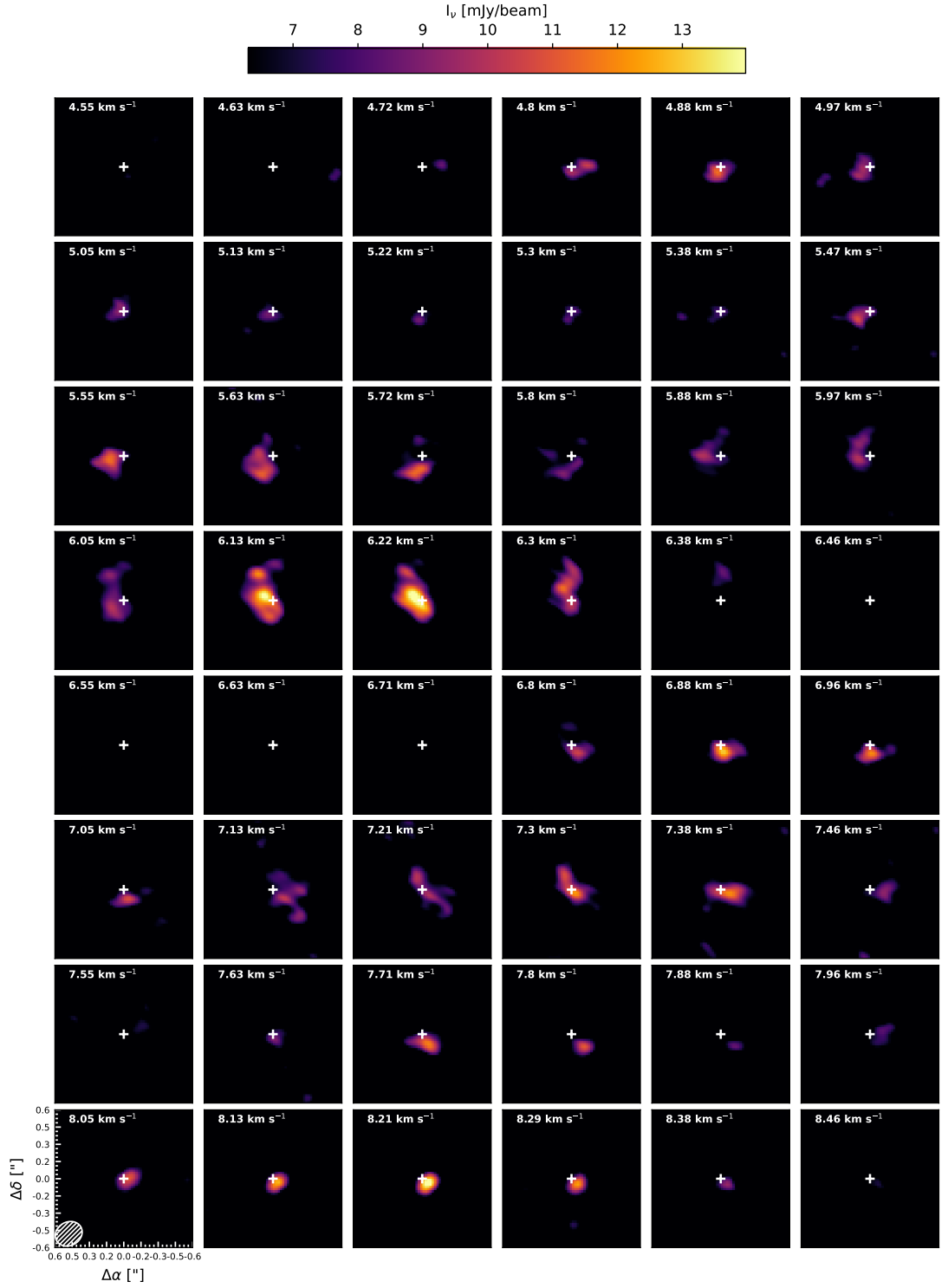


FIGURE B.3: Channel maps for SO 6_5-5_4 emission in IRAM04191+1522 from $V_{LSR} = 4.55 \text{ km s}^{-1}$ to 8.46 km s^{-1} . The emission shown is clipped at $3\sigma_{\text{rms}}$ and above ($\sigma_{\text{rms}} = 2.102 \text{ mJy/beam}$). The beam ($0''.24 \times 0''.20$ at 119.4°) is shown in the bottom left of the lower left channel map. The systemic velocity is 6.5 km s^{-1} . The source position is indicated by the white cross.

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