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A Study of Herbig Ae Be Star HD 163296: Variability in CO and OH

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A Study of Herbig Ae Be Star HD 163296
Variability in CO and OH

A Thesis
Presented to
the Graduate School of
Clemson University

In Partial Fulfillment
of the Requirements for the Degree
Master of Science
Physics and Astronomy

by
Diana Lyn Hubis Yoder
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Abstract

Spectra of the Herbig Ae Be star HD 163296 show variability in the CO and OH ro-vibrational emission lines. Documented here is the variance of OH emission lines between measurements separated by a period of three years. By comparing this variability to the variability of CO, we aim to determine whether the two are linked through an event such as disk winds. We find the profiles of OH and CO taken within three months of each other to have exceedingly similar profile shapes. However, the variability seen between the OH data sets needs further work to verify whether the changes we see are reproducible.
Dedication

I would like to dedicate this thesis to my husband, Marcus. Without whom, I don’t know where my path would have taken me.
Acknowledgments

Over the past two years, I have been confronted with a multitude of exciting, new challenges. Most of which, I could not have faced without the support of my friends and family here in Clemson. From the stacks of chores my husband helped with, to the late night study (or ’de-stress’) sessions with classmates, the support and love have made these years wonderful.

I would also like to thank my parents and older sister for raising me to believe that I can take on any challenge I choose. The world is full of opportunities, I need only decide which ones I wish to pursue.
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Chapter 1

Introduction to Herbig Ae Be Stars

Herbig Ae Be stars (HAeBes) are intermediate mass, pre-main sequence stars which have strong emission lines in the near-infrared (NIR). These young stars range between 2-10 solar masses and have a spectral type of A or B indicated by their temperature (Herbig, 1960). HAeBes will evolve to become main sequence A type stars with the potential to have planets formed from a circumstellar disk which revolves around them (Su et al., 2006). This chapter describes how HAeBes are formed and why one such star in particular, HD 163296, is of importance.

1.1 Molecular Cloud Collapse

The gas and dust which composes the interstellar medium (ISM) consists of approximately 70% hydrogen by mass. Most of the rest consists of helium with a small mix of other metals (Wilson & Rood, 1994). The hydrogen is made up of neutral hydrogen in the ground state (H I), ionized hydrogen (H II), and molecular hydrogen, (H$_2$). These gases and dust in the ISM occasionally group together, forming higher density regions called molecular clouds.
An example of one such molecular cloud is the Orion Nebula in Figure 1.1. The Orion Nebula is one piece of a larger system called the Orion Molecular Cloud Complex. It contains approximately 1000 solar masses while being only 9 light years in diameter (Zuckerman, 1973). This molecular cloud is a stellar nursery where many stars are being formed.

Figure 1.1: The Orion Nebula. Image credit goes to ESA/Hubble
1.1.1 Jeans’s Mass

Protostar formation begins with small deviations from hydrostatic equilibrium in the molecular clouds. A simplified version of the mass necessary for collapse can be found with the following derivation. Using the virial theorem which states the average energy is equal to half of the potential energy, \( E = \frac{1}{2} U \), and the fact that energy is equal to the sum of the kinetic energy and potential energy gives us the following equation:

\[
2K = -U
\]

The motion of \( N \) atoms inside the cloud creates kinetic energy equal to \( K = \frac{3}{2} NkT \), while the gravitational potential of the entire cloud is \( U = -\frac{3}{5} \frac{GM^2}{R} \). Combining these equations gives us the condition for virial equilibrium.

\[
3NkT = \frac{3 \ G M^2}{5 \ R}
\]

If the left hand side of the equation was greater, then thermal energy would cause expansion of the cloud. However, if the right hand side of the equation was greater, collapse would begin.

Looking at the latter scenario and assuming a spherical cloud with constant density, we find \( R = \left( \frac{3M}{4\pi\rho} \right)^{1/3} \). We can also replace the number of atoms, \( N \), with the mass of the cloud divided by the mass per particle. This gives

\[
M > \left( \frac{5kT}{Gm} \right)^{3/2} \left( \frac{3}{4\pi\rho} \right)^{1/2} = M_J
\]

This defines the Jean’s Mass, \( M_J \). The mass of the molecular cloud must be larger than the Jean’s Mass before collapse can occur (\( M_C > M_J \)). In actual molecular clouds, the shape and dynamics are not this simple. Nebula such as the one if Figure
1.1 have varying shapes and densities, making it more difficult to determine the necessary mass for collapse. However, this simplified result tells us enough to understand how collapse of a star begins (Carroll & Ostlie, 2007).

### 1.1.2 Free Fall Time

If the density of the cloud is uniform throughout, it will take the same time for all of the cloud to collapse with the free fall time, $t_{ff}$, which depends only on density. This is called Homologous Collapse. If the cloud is not uniform, $t_{ff}$ will be shorter for the center of the cloud and it will collapse from the inside-out. The following is a derivation of the free fall time for the Homologous case.

We begin with the local acceleration of gravity a distance $r$ from the center of a spherical cloud.

$$\frac{d^2r}{dt^2} = - \frac{GM_r}{r^2}$$

By replacing $M_r$ (the mass to distance $r$) with the volume and initial density of this sphere, and multiplying both sides of the equation by the velocity of the surface of the sphere, we obtain

$$\frac{dr}{dt} \frac{d^2r}{dt^2} = - \frac{4\pi}{3} G \rho_0 r_0^3 \frac{1}{r^2} \frac{dr}{dt}$$

Integrating this with respect to time gives

$$\frac{1}{2} \left( \frac{dr}{dt} \right)^2 = \frac{4\pi}{3} G \rho_0 r_0^3 \frac{1}{r} + C_1$$

Assuming the initial velocity of the sphere’s surface to be zero, we find $C_1$ to be

$$C_1 = - \frac{4\pi}{3} G \rho_0 r_0^2$$
We can now solve for the velocity at the surface of the sphere.

\[ \frac{dr}{dt} = -\left[ \frac{8\pi}{3} G \rho_0 r_0^2 \left( \frac{r_0}{r} - 1 \right) \right]^{1/2} \]

Now, in order to obtain an expression for the position of the cloud as a function of time, we have to make a series of substitutions. Using

\[ \theta = \frac{r}{r_0}, \quad \chi = \left( \frac{8\pi}{3} G \rho_0 \right)^{1/2}, \text{and} \quad \theta = \cos^2 \xi \]

we can find the equation

\[ \frac{d\theta}{dt} = -\chi \left( \frac{1}{\theta} - 1 \right)^{1/2} \]

which becomes

\[ \cos^2 \xi \frac{d\xi}{dt} = \frac{\xi}{2} \]

We can now integrate with respect to \( t \) to find

\[ \frac{\xi}{2} + \frac{1}{4} \sin 2\xi = \frac{\chi}{2} t + C_2 \]

\( C_2 \) can be found by setting \( r = r_0 \) when \( t = 0 \) so that \( \theta = 1 \) and \( \xi = 0 \). This makes \( C_2 = 0 \).

We now have an equation which describes the motion of the cloud under gravitational collapse.

\[ \xi + \frac{1}{2} \sin 2\xi = \chi t \]

Using this, we can determine the free fall timescale by saying \( t = t_{ff} \) when the radius
of the sphere is zero, or \( \theta = 0 \) and so \( \xi = \pi / 2 \). Once we substitute \( \chi \) back in, this gives

\[
t_{ff} = \left( \frac{3\pi}{32G\rho_0} \right)^{1/2}
\]

We know now how long it would take for a spherical cloud of equal density to collapse into a star (Hayashi, 1966). Unfortunately, molecular clouds do not behave quite this uniformly, but this is a close enough approximation for most purposes (Carroll & Ostlie, 2007).

### 1.1.3 Initial Mass Function

If the cloud greatly exceeds the Jean’s Mass, then a process called fragmentation will occur where sections of the cloud will start to collapse individually. This can cause a cascade effect of the largest stars living and dying spectacularly before the smaller stars have finished accreting matter, resulting in heavy metals being dispersed throughout the region of space.

The initial mass function (IMF) is the theoretical distribution of high to low mass stars which should be created when a molecular cloud fragments, based on the initial mass of the molecular cloud. An IMF can be determined for a stellar cluster by studying the luminosity of the stars, but it is still unsure whether a universal IMF can be found. In general, the IMF is modeled as a power-law fit of the form

\[
\xi(M_*) = C \left( \frac{M_*}{M_\odot} \right)^{-(1+x)}
\]

where \( C \) is a normalization constant and \( x \) depends on the relevant mass range (Salpeter, 1955).
1.1.4 Zero-Age Main Sequence

The Hertzsprung-Russell (HR) diagram is the chart which shows stars’ relationships between temperature, luminosity, and mass. The main sequence is the region of the HR diagram where the stars have hydrogen fusion occurring in their cores. Most stars spend the majority of their lives here (See Figure 1.2). Zero-age main sequence (ZAMS) is when a star first reaches the main sequence and first begins equilibrium hydrogen burning. This marks the beginning of ’adulthood’ for the star. The time it takes for a protostar to reach this point is inversely related to its mass. Stars with a fraction of a solar mass take millions of years to reach ZAMS, while a 60 solar mass star takes only a few thousand years (Hayashi, 1961).

Figure 1.2: Diagram showing the relationship between various stellar parameters. Image credit goes to ESO.
Pre-main sequence stars experience rapid changes in temperature and luminosity before finding their place on the main sequence. Figure 1.3 shows the path protostars of various masses take before landing on the main sequence (Bernasconi & Maeder, 1996). Lower mass stars will take the nearly vertical routes in Figure 1.3 called the Hyashi Track. HAeBe stars spend very little of their time on the Hyashi track, instead much of their young life lies on the nearly horizontal route; increasing temperature while keeping the same luminosity (Hayashi, 1961).

Figure 1.3: Pre-main sequence life of various mass stars. The dotted line represents the birthline and the various mass stars follow their tracks to the left until they come to an end at ZAMS. The nearly vertical portion of each line represents the Hyashi Track. The approximate time scale for a protostar traveling these tracks is given by the point listed in the bottom left (Palla & Stahler, 1994).
1.2 Stellar Evolution

We have now followed a protostar from birth in a molecular cloud to adulthood as a member of the main sequence. Now, we need to look back at a few of the dynamic processes associated with these transitions. Stars are constantly moving and changing, resulting in violent outbursts and an active environment. This causes the material surrounding them to sometimes behave erratically and it is important to understand the possible effects on the region.

1.2.1 Convection

Convection is caused by parcels of material in the star rising through the hot plasma to areas much cooler than the ones close to the core, where fusion is taking place. These parcels rise up slightly and are much warmer than the surrounding material, causing them to have a lower density. Now, the parcels continues to rise up due to its newfound buoyancy. This causes other parcels of plasma to now be more dense than the surrounding area, and these sink down towards the core. This adiabatic cycle is known as convection, and keeps the interior of the star uniform with constant mixing of the material.

Convection is not always present throughout a star. In many situations, the parcel of material will exchange heat with its surroundings instead of changing its buoyancy. The possibility of convection is marked by the Schwarzschild criterion which states whether the material will respond adiabatically to its surroundings.

\[-\frac{dT}{dZ} < \frac{g}{C_p}\]

Where \(Z\) is the height from the center of the star and \(C_p\) is the heat capacity at
constant pressure for the star. If this relationship is not true, then the material is considered to be unstable against convection and a convection current will be created. In regions of a star where the plasma has been ionized, convective stability is more difficult to satisfy due to the adiabatic exponent being lowered and the high opacity. This effect is especially apparent in the outer regions of the star where hydrogen and helium are cool enough to be partially ionized (Carroll & Ostlie, 2007).

1.2.2 Stellar Dynamo

Convection causes ionized plasma to be constantly moving through the star, creating a turbulent environment. The star, also has a rotation gradient between the poles and equator, with the plasma around the middle moving much slower than across the poles. These two phenomena work together to create strong magnetic fields across the star. This dipole is thought to be responsible for several different mechanisms surrounding the star, including outflows and accretion. The strength and consistency of the dipole is dependent on both the age and mass of the star. We will discuss each of these variables in the next two sections.

1.2.3 Mass Differences between Protostars

The mass of a protostar is an important variable during development and evolution of a star. T Tauri Stars are the low mass, pre-main sequence cousins to HAeBes. They range from 0.5-2 solar masses and are known for their large, rapid, irregular variations in luminosity. These variations occur rapidly over the course of several days. T Tauri stars are also known for their strong emission lines in the Balmer series of hydrogen, Calcium II, Iron, and Lithium wavelengths. Emission in the forbidden lines of [O I] and [S II] are also present, which indicate an extremely
low gas density. Another aspect of T Tauri stars is the unique shape of the emission lines, which indicates significant mass loss. The shape of the emission line is shifted relative to the absorption line.

T Tauri stars have stronger magnetic fields than the HAeBes and maintain the field for longer after reaching the main sequence. The cause for this difference is not fully understood, but is believed to be linked to the convective envelopes of the low mass stars. In comparison, higher mass stars have a convective core, with a radiative envelope.

1.2.4 Magnetic Field Over Time

As the protostar ages, recent studies show that the magnetic field of different mass stars behave differently over time (Hubrig, 2009). T Tauri stars appear to maintain strong magnetic fields throughout their pre-main sequence lives and then continue to have strong magnetic fields once they have reached adulthood. The behavior of HAeBes’ magnetic fields over time is a bit more variable. The internal structure of a HAeBe star changes drastically during its evolution towards ZAMS. Figure 4.3 shows the life cycle of a pre-main sequence star as its structure changes between convective and radiative.

Lower mass HAeBes start their lives as fully convective, quickly developing a radiative core, and then spending much of their remaining pre-main sequence life as fully radiative. Just before reaching ZAMS, they develop a small convective core. Higher mass HAeBe stars can even begin life as fully radiative, only developing any convection as they approach ZAMS.

In contrast, all T Tauri stars begin life as fully convective. They eventually develop a radiative core, but retain their thick convective envelope. This convection
gives them their strong magnetic field throughout life. HAeBes do not have this extensive convective zone, and hens develop much weaker fields (Neiner et al., 2015). Hubrig (2009) found that the magnetic field of HAeBes is strongest when they are very young and weakens as they approach ZAMS. The decline of this field causes changes in the accretion and outflow rates which will be discussed further in Section 1.3.5.

Figure 1.4: Diagram showing how the stellar interior changes during a protostar’s approach to ZAMS. Different mass stars will begin with different interiors, but follow a similar pattern as they approach adulthood (Neiner et al., 2015).

1.3 Protoplanetary Disks

Both types of pre-main sequence stars are enveloped in a ring of dust and gas, which we call the protoplanetary disk. This disk is responsible for several de-
velopmental aspects of the young protostar. First, we will look at how it is formed from conservation of momentum and discuss the potential for planetismals. Then, we will observe how the disk regulates the angular momentum of its parent star and the outflows this may create.

1.3.1 Conservation of Angular Momentum

Protostars go through several distinct phases during formation and we label these phases with classes. These classes and their most distinctive attributes can be seen in Figure 1.5. Class 0 is when the nebula is still roughly spherical and most of the stellar accretion is taking place. This stage lasts about $10^4$ years. At an age of $10^5$ years, Class I begins and a solid core has been created from accretion while the rest of the material in the cloud is beginning to form a disk around it. This gas and dust make up the circumstellar disk.

Young stars have circumstellar disks due to conservation of angular momentum. In the absence of external torques, the angular velocity increases as material begins to gather at the center of the system. However, if this were the whole picture, the disks would eventually be spinning at such high speeds that they would rip apart. Instead, some angular momentum must be transferred away. One way this occurs is through magnetic fields anchored to convection zones in the protostar and coupled to ionized stellar winds. This will slow rotation by applying a torque against the disk.

This Class is indicated by a positive spectral energy distribution slope (SED). SED’s are plots of the flux density of a star versus the wavelength of observed light. The average SED tells us much about a star’s composition and phase of life. Class II starting at $10^6$ years old, marks the start of contraction of the star towards the main sequence on the HR diagram.
Figure 1.5: The various Classes of young stellar objects and how their disks take shape compared to the SED given on the left. (Bachiller, 1996)
Once the protostar begins Class II, most of the star’s mass has now been collected and the disk is optically thick due to the remaining debris, causing the SED slope to be flat. At $10^7$ years old, Class III begins and is marked by a negative SED slope. The disk is now mostly empty because of accretion and the outflow material, and hence optically thin.

Between Class II and III, there is a very interesting period where the inner parts of the disk will become optically thin as the dust disappears (to around 70 AU), while the outer regions are still optically thick. These are called transition disks and are objects of much interest in the scientific community. The age of a transition disk coincides with ideal planet formation conditions and by studying the dips in SED, we can determine whether or not there are gaps in the disk. These gaps are several AU wide and will result in the inner disk being depleted of material and ending the star’s accretion of matter.

1.3.2 Disk Shape

The shape of an accretion disk is of much importance, and there are several theories around what geometry a given disk has. Current observations show that HAeBe stars have a gap between the star and the inner rim of the disk. This inner rim is heated directly by the star and consequently becomes much hotter than the rest of the disk and develops a puffed up region. The rest of the disk can either be flared or self-shadowed. A flared disk gains height further from the star, while a self-shadowed disk remains close to the solar plane Meeus et al. (2001). These shapes are shown in Figure 1.6.
In order to determine the structure of the disk, the vertical force due to the gravity of the star must balance with the vertical pressure of the disk. Two core assumptions are made in order to simplify the calculations: (1) The mass of the disk must be less than or equal to 0.1 solar mass so that we can neglect the gravity of the disk, and (2) That there is sufficient cooling along the surface area of the disk to allow low disk temperatures and pressures. Both assumptions prove to be reasonable for accretion disks.

The gravitational force of the star depends on the cylindrical radius $r$ at a height $z$ around a star of mass $M_\ast$. The vertical component is given as

$$g_z = \frac{GM_\ast}{(r^2 + z^2)^{3/2}}$$

This must balance with the vertical pressure gradient in the gas $(1/\rho)(dP/dz)$, where the pressure can be written as $P = \rho c_s^2$ with $c_s$ being the speed of sound. Equating
the acceleration due to pressure to that of gravity gives

\[ \frac{c_s^2 d\rho}{c_s dz} = \left[ \frac{GM_* z}{(r^2 + z^2)^{3/2}} \right] \]

Solving this for a thin disk \( z \ll r \) and \( g_z \approx \Omega^2 z \), where \( \Omega = \sqrt{GM_*/r^3} \) is the Keplerian angular velocity, gives

\[ \rho = \rho_0 e^{-z^2/2h^2} \]

With the mid-plane density \( \rho_0 \) written in terms of the full-plan surface density \( \Sigma \),

\[ \rho_0 = \frac{1}{\sqrt{2\pi}} \frac{\Sigma}{h} \]

and \( h \) is the vertical disk scale-height given as \( h = \frac{c_s}{\Omega} \). If we parameterize the speed of sound by \( c_s \propto r^{-\beta} \), then the aspect ratio of the disk varies as

\[ \frac{h}{r} \propto r^{-\beta+1/2} \]

This shows that the disk will flare upwards as the radius increases (Pringle, 1981). This derivation requires that \( T(r) \propto r^{-1} \), which proves to be reasonable.

The flared disks are considered Group I HAeBes while the self-shadowed disks (with different temperature profiles) are Group II (Meeus et al., 2001). However, van Boekel et al. (2005) discovered a way to distinguish the two groups through a protostar’s luminosity and color at various wavelengths. Group I and II YSO’s are separated by their IRAS \( m_{12} - m_{60} \) color versus \( L_{NIR}/L_{IR} \) values. A diagram of a group of 24 isolated HAeBes and their ratios of luminosity to color is shown if Figure 1.7. Group I stars have the relationship \( L_{NIR}/L_{IR} \leq (m_{12} - m_{60}) + 1.5 \) while group II stars have \( L_{NIR}/L_{IR} > (m_{12} - m_{60}) + 1.5 \) (van Boekel et al., 2005).
Figure 1.7: Classification of HAeBes based off their NIR to IR luminosity ratio and IRAS $m_{12} - m_{60}$ color (defined as $m_{12} - m_{60} = -2.5\log F_{12}/F_{60}$) (van Boekel et al., 2005). HD 163296 is represented as diamond 22.

1.3.3 Line Profiles

HAeBe stars frequently show CO line emission in their spectra. This emission shows details of the inner regions of the disk where planet formation occurs. These lines are very important because the shape reveals the geometry of the emitting gas. However, we have to first separate out observational and instrumental effects before we can discern any valuable information.

One way the line might become distorted is by having a low signal to noise ratio. This can result in line profiles that are asymmetric, have double peaks, or have
flatter tops than they are supposed to. Low spectral resolution can also interfere with a line profile by smoothing out a double peak or other features of the line. Finally, slits might lose part of the velocity field, resulting in inaccurate geometry.

If instrumentation has gone correctly, we can expect a CO line emission of a highly inclined disk to have a double peak (Smak, 1981). The Keplerian rotation causes the inner regions of the disk to be rotating faster than the outer regions. Hence, the wings of a line profile are determined from the speeds of the inner disk, while the peaks are defined by the outer disk. Once Doppler shift has been accounted for, the center of the emission should have no velocity. These relationships are shown in Figure 1.8. Scientific reasons for there not to be a double peak include a low inclination of the disc, a composite emission from both the inner and outer regions of the disk, and disk winds or outflows.

Figure 1.8: The Figure on the left represents the observed speeds of different regions of the disk as observed from edge-on. The Figure on the right shows the resulting line profile shape. The more gas rotating at a specific speed, the higher the intensity of the line profile (Beckwith & Sargent, 1993).
1.3.4 Planet Formation

Gaps present in the disk indicate there might be large objects in orbit around the system (Najita et al., 2007). These objects could be young planets. Planet formation is still an area of much debate. There are several models in place that could explain how planets form, but they all have their own obstacles. The generally accepted theory is that dust in the accretion disk begins to collect together, forming larger and larger pieces until they begin to have their own gravitational pull. At this point, the object is able to clear out its orbit by collecting all nearby dust, and becomes a planetesimal. This planetesimal will either continue to grow and become a planet, or will be smashed apart by other objects in the disk.

There are several methods to find these gaps and identify possible ongoing planet formation. One way is to look for emission lines within an AU of the protostar. CO’s high dissociation temperature makes it the perfect candidate to trace the shape of the circumstellar disk. By observing its presence and shape, gaps in the disk can be documented (Najita et al., 2007).

Another way to identify planet formation is through the eccentricity of the disk (Regály et al., 2010). The presence of a large companion causes a distinctive asymmetric line profile in the spectrum of the disk. Asymmetry is caused by the relative speeds of the rotating gas. Gas near the periastron is moving faster than the apastron and so a doppler shift is present within the line profile. A model of a line profile with various eccentricities is shown in Figure 1.9.
1.3.5 Angular Momentum Regulation

We have been discussing many unique properties of disks as though they are separate bodies. However, the circumstellar disk and parent star are closely linked and changes in one will greatly affect the other during these early stages of development. The star’s magnetic field extends into the accretion disks and facilitates the transfer of material from the disk to the star. The inflowing gas and dust is then braked by the magnetic field until its angular velocity stabilizes with the corotation radius of the star. This link is shown in Figure 1.10. More importantly for HAeBes, is the link between disk and star which regulates angular momentum through outflows. Dispersing material away from the system helps to slow the rotation of the star and disk which helps to maintain equilibrium. Outflow through disk winds and Herbig Haro jets are discussed in the next section.
1.3.6 Outflow Away from the Star

There are several methods for a star to disperse excess angular momentum by expelling material away from it. The most general of which is via a disk or stellar wind. These winds are found around both T Tauri and HAeBes and can make observations of the star difficult because of the variable nature of the wind. The cause and affect of disk winds can vary between observations taken a few years apart, and is a current obstacle to determining the standard makeup of young stellar objects (YSO). Figure 1.10 shows how the wind blows material away from the plane of the disk along lines created by the protostar’s magnetic field.

Herbig-Haro Objects [HH] are the result of narrow beams of gas that are ejected in opposite directions of the star roughly perpendicular to the disk. As the outflow supersonically collides with gas and dust near the disk and previous ejecta,
all of this material then becomes shockheated nebulae, the HH object.

HH jets are commonly associated with T Tauri stars, but are more recently associated with HAeBes as well (Grady et al., 2000). As they excite gas, emission lines are created in the spectrum which indicate composition and direction of the jet. Each jet evolves over the course of a few decades and new jets can form by the time the last dissipates.

Figure 1.11 shows one such HH jet being ejected from the poles of a new born star in the Orion Molecular Cloud Complex. As the star ejects material, supersonic shock fronts are created and heat the surrounding gas enough to form this beautiful display in the infrared.

1.4 HD 163296

HD 163296 is an isolated, pre main sequence star of spectral type A3Ve (Grady et al., 1998) located 118 ± 11 pc from our solar system (van Leeuwen, 2007). It has a mass of 2.3$M_\odot$ (Montesinos et al., 2009), an inclination of 46° ± 4°, and a position angle of 128° ± 4° (Isella et al., 2007). Montesinos et al. (2009) estimates the age of HD 163296 to be 5±0.3−0.6 Myr. Meeus et al. (2001) observes that the disk is flat while van Boekel et al. (2005) finds its luminosity ratio to be greater than the IRAS color, and therefore both methods find it to be a Group II, Herbig Ae star.

This star has been a topic of much interest for many years. Numerous observations in a variety of wavelengths have been made, making it a good candidate for observing how a classical Herbig Ae star changes over relatively short time periods. Of specific interest are molecules such as CO and OH which can act as tracers for H$_2$. The existence of consistent data means that we can compare observations of these molecules at various time periods over the last two decades.
Figure 1.11: A young protostar hides behind a cloud of gas and dust while Herbig Haro 24 illuminates the ejected material in the Orion Molecular Cloud Complex. Image credit goes to ESA/Hubble.
1.4.1 Herbig Haro Jet 409

In 2000, Grady et al. discovered that T Tauri stars are not the only pre main sequence stars to produce Herbig Haro Objects. Coronagraphic imaging of HD 163296 showed evidence of several knots of nebulosity emanating from the star, roughly perpendicular to the plane of the disk, in addition to an oval region of gas centered around the star. The ionization of the knots shows that they are low-excitation Herbig Haro knots (Grady et al., 2000). These knots (HH 409) are the remains of outflow created at 16 year intervals dating back to the 1800’s (Ellerbroek et al., 2014). Figure 1.12 shows an image of HD 163296 along with a schematic of HH 409.

Figure 1.12: Coronographic image of HD 163296 taken with STIS, next to a long-slit spectrum, and a schematic of the placement of HH 409. (Devine, D. , Grady, 2000)
1.4.2 Importance

As an isolated HAeBe star, HD 163296 bridges the gap between classical pre-main sequence and main sequence stars, because although it is of spectral type Ae and exhibits infrared excess and emission lines, it is not associated with dark clouds or reflection nebulae like most HAeBes. This absence of nearby material makes it an important object to study in order to understand the formation and evolution of HAeBe systems. However, in order to apply this knowledge to other systems, we must first understand what makes HD 163296 so variable. In the following chapter, we will review the previous work surrounding HD 163296 and its variable emission in the NIR.
Chapter 2

HD 163296 Variability

The most notable aspect of HD 163296 is how much variability is found in the spectrum over the years. The photometric variability and NIR excess fluctuations make HD 163296 a rapidly evolving system, and measuring these variations tells us what might be going on at the inner edges of the disk. Optical observations date back to 1980 and are taken with good frequency up until the most recent in 2012. The NIR bands were first observed in 1979, but were not studied in detail until the early 2000s (Sitko et al., 2008). In this Chapter, we will review the work already done with the changes in flux for the last decade, the latest results surrounding HH 409, and give a detailed discussion of the variability seen in CO.

2.1 Infrared Variability

Over the last 35 years, flux measurements for HD 163296 have been taken in nearly every wavelength between 0.1 and 10000 \( \mu m \). Sitko et al. (2008) has compiled all of these measurements into a single plot shown in Figure 2.1. This plot helps to show why HD 163296 is considered a Class II, Group 2 HAeBes. As you can see, the
flux in units of $\lambda F_\lambda W m^{-2}$ is only slightly lower at $60\mu m$ than at $8\mu m$, making the SED mostly flat here and therefore a Class II YSO.

Figure 2.1: The spectral energy distribution of HD 163296. This is a compilation of many works and various instruments (Sitko et al., 2008).

Taking a more refined look at the data between 1 and 15 $\mu m$, we can see where the flux appears to deviate the most. Figure 2.2 shows the flux in this range for data taken between 1979 and 2006. Between 2002 and 2005, prominent fluctuations in the NIR at 1-5 $\mu m$ are observed. This change in flux is commonly attributed to structural deviations along the inner rim near the dust sublimation zone (DSZ), resulting in the inner rim’s distance from the star varying over time. Another explanation of the fluctuations in the NIR could be the removal of dust from the inner rim. An event such as a disk wind could be driving gas and dust outward, away from the disk (Sitko et al., 2008).
Figure 2.2: The flux density of HD 163296 from 1-15\(\mu\)m after the stellar continuum has been removed. This is a compilation of many works and various instruments. Adopted from Sitko et al. (2008).

The fluctuations seen between 2002 and 2005 represent two different epochs in the star’s life-time, indicating a change from the normal flux state. Sitko et al. (2008) ran models of the disk structure of HD 163296, trying to determine how the DSZ might be changing over the different epochs. They found the DSZ to be at 0.29 AU for the low flux state of the star, and 0.35 AU for the high flux state. However, research done by Vinković et al. (2006) shows that the presence of silicate features at 10 \(\mu\)m indicate that a vertical inner edge cannot explain the NIR emission. If silicates are present, the dust will not have a perfectly gray opacity. Sitko et al. (2008) did discover significant silicate fluctuation in his data, indicating that changes in the DSZ are not responsible for the flux.
Another explanation for the outburst in 2002 is an increased illumination of the star. However, comparing the spectra of 2002 and 2005 shows that the flux was constant at 0.8 µm. This steady flux in the NIR proves the outburst cannot be caused by variable heating by the star in the optical (Sitko et al., 2008).

The possibility of heating from accretion shock created by magnetospheric accretion remains. Gas and dust follows the field lines of the stars’ magnetic dipole, until it falls onto the star emitting shocks in the X-ray. These X-rays are absorbed by the stellar atmosphere and re-emitted as UV radiation. HD 163296 is known to possess hot accreting gas. This, along with the highly collimated outflows, indicate magnetospheric accretion processes are likely present (Sitko et al., 2008).

Shu et al. (1994) models the X-Wind created by the ionization of the gaseous disk by accretional heating and high energy stellar and accretion shock photons. This model discusses the production and collimation of outflowing jets and considers the possibility of reconnection events in the magnetic field altering the outflow rate and structure of the disk near DSZ. Sitko et al. (2008) considers the possibility for X-Winds in HD 163296 and determines that the time scales for these changes are similar to the time scales of Herbig Haro objects.

### 2.2 Herbig Haro 409

There are ten distinct knots of nebulosity emanating from the HD 163296; HH 409 A, A2, A3, B, B2, C-G. Figure 2.3 shows these knots in several different emission lines. The original three knots; A, B, and C, were notable because of their associated Lyα emission; this was the first detection of a Lyα bright jet driven by a Herbig Ae star Grady et al. (2000).
Devine, D. , Grady (2000) discusses the emission of the spectrum and how the blueshifted region extends toward the southwest while the redshifted follows the northeast axis, proving that the source is emanating from the star perpendicular to the disk. Later, Ellerbroek et al. (2014) determined that the jet was perpendicular to within 4.5° and that there is a distinct asymmetry between the lobes above and below the disk. The blue lobe is seen to be travelling faster (220-300 km s$^{-1}$) than the red lobe (130-200 km s$^{-1}$) (See Figure 2.3). This, along with the higher ionization fraction and electron temperature in the blue lobe, causes the blue lobe to have higher excitation conditions.

The asymmetry between the lobes might be caused by an asymmetric disk structure or a magnetic field, but the equal energy input and dual outflows indicate that there might be a single driving mechanism controlling the jet launches. Interferometric data of HD 163296’s disk can be seen in Figure 2.4. Here, you can see where the blue jet lobe is thought to be.
Ellerbroek et al. (2014) compiled all of the data for HH 409 knots into a plot of the distance a knot is from the source versus the year the data was taken. This allowed him to use the trajectories of the knots to estimate their launch epochs. These epochs represent the intervals during which the bright knots formed inside the jet, close to their source. The position of the knots based off of the FWHM of the spatial profiles and the projected launch epochs are shown in Figure 2.5.
Based off of these projections, the bright shock fronts appear simultaneously for each lobe and with a period of 16.0±0.7 yr. By combining these newfound launch epochs with the fluxes seen in HD 163296 over the last 70 years, Ellerbroek et al. (2014) was able to show that the increases in flux coincides with the launch of the knots. This relationship is shown in Figure 2.6.
The jet causing the knots to appear, consists of two distinct regions, with the various knots embedded inside. The region nearer the star contains a narrow, high-velocity jet, while the outer region is a lower-velocity disk wind. This disk wind contains the visible dust clouds as they interact with gas and dust around the system. These dust clouds are most likely responsible for observed photometric variability in the star. Once the dust cloud is lifted high enough above the plane, it will be exposed to direct stellar light and will warm up, emitting thermal radiation in the NIR. Figure 2.7 shows a schematic of the placement of the high and low velocity winds, with a dust cloud for reference.
The thermal radiation from the lifted dust cloud will cause variability in measurements of the spectrum in the NIR. A disk wind of appreciable force is required to lift this cloud above the disk. An alternative source of lifting is the existence of a companion on a wide eccentric orbit causing tidal disruption and forcing the dust up and out of the disk (Ellerbroek et al., 2014).

A possible result of HH 409 is that the disk clearing act of the outflow increases the in fall rate onto the star. This would rejuvenate the outflow potential, causing subsequent expulsions. However, there is no evidence of a relation between the accretion rates and the outflow (Ellerbroek et al., 2014).
2.3 CO Variability

CO emission from HD 163296 is known to be variable (Bertelsen, 2014). NIRSPEC data collected in 2001 (Salyk et al. 2011; S11) and 2002 (Brittain et al. 2007; B07) and CRIRES spectra collected with the European Space Observatory’s (ESO) Very Large Telescope (VLT) from 2012 (Bertelsen 2014; B14) show this. Analysis of these data sets showed a continuum flux at 4.77 $\mu$m to be 16.3 Jy for B14 and 12.1 Jy for S11. Each of these values have a 10-20% uncertainty due to the differences in width of the point spread function (PSF). All three of these data sets showed wide emission lines, indicating that the CO ro-vibrational lines are emitted within an outer radius of 2-3 AU of the star (Bertelsen, 2014).

The shape of the line profiles vary between each data set. Each of the relevant line profiles are seen in Figure 2.8, while an averaged comparison of each is shown in Figure 2.9. B14 showed single peaked line profiles for low and mid J values, but flat topped asymmetric lines with wide shoulders for high J values. S11 found the same single peaked profiles for low J values, but the mid and high values showed double peaked or flat topped profiles. Overall, the S11 data had wider FWHM than the B14 data. The spectra from B07 show a mix between single and double peaked emission lines, with no obvious relation between the high or low J (Bertelsen, 2014).

The CRIRES instrument has a higher spectral resolution than NIRSPEC, and should therefore be able to discern any possible double peak. However, despite the superior signal to noise ratio, the CRIRES data set shows no hint of a double peak. The decade long gap between the data sets leaves room for variability in the star to cause these differences.
Figure 2.8: Line profiles for the CO data taken from Bertelsen (2014), Salyk et al. (2011), and Brittain et al. (2007) in that order. The compilation comes from Bertelsen (2014).
Figure 2.9: Mean line profile comparisons for the CO data taken from Brittain et al. (2007), Salyk et al. (2011), and Bertelsen (2014). Here, we compare the profiles for various values of J separately as was all as the averaged profiles (Bertelsen, 2014).

The first possibility for these differences in line shape, lies in the NIR brightness increases already known about for HD 163296. Three epochs during 1986, 2001-2002, and 2011-2012 have been shown to increase NIR brightness by 10%. However, seeing as all three data sets were taking during these epochs, some changes in flux are too be expected. Hence, this increase in brightness does not fully explain the different shapes in the line profiles (Bertelsen, 2014).

Additionally, Table 2.1 shows that while the FWHM values are not consistent between datasets, the FWZI (Full Width at Zero Intensity) does agree for B14 and S11. This indicates that the width between profiles remains constant, while the intensity and shape vary.
Table 2.1: FWHM from a Gaussian fit and FWZI from manual measurements by Bertelsen (2014). All numbers are given in km/s

<table>
<thead>
<tr>
<th></th>
<th>B14</th>
<th>S11</th>
<th>B7</th>
</tr>
</thead>
<tbody>
<tr>
<td>FWHM, all J</td>
<td>55.0±2.5</td>
<td>73.9±1.4</td>
<td>60.0±1.4</td>
</tr>
<tr>
<td>FWHM, J&lt;10</td>
<td>54.5±3.9</td>
<td>66.9±1.8</td>
<td>45.8±1.6</td>
</tr>
<tr>
<td>FWHM, 10&lt;J&lt;20</td>
<td>50.9±4.9</td>
<td>74.3±3.8</td>
<td>—</td>
</tr>
<tr>
<td>FWHM, J&gt;20</td>
<td>59.2±4.3</td>
<td>85.0±2.8</td>
<td>78.9±2.2</td>
</tr>
<tr>
<td>FWZI, all J</td>
<td>130±10</td>
<td>145±10</td>
<td>130±10</td>
</tr>
<tr>
<td>FWZI, J&lt;10</td>
<td>132±7</td>
<td>140±10</td>
<td>110±20</td>
</tr>
<tr>
<td>FWZI, 10&lt;J&lt;20</td>
<td>125±15</td>
<td>130±15</td>
<td>—</td>
</tr>
<tr>
<td>FWZI, J&gt;20</td>
<td>128±7</td>
<td>145±15</td>
<td>150±10</td>
</tr>
</tbody>
</table>

Bertelsen proposes two alternate theories as to why the line profiles differ over the years. The first suggests that the low J values which show single peaks are in fact composites of multiple emitting regions. One component is due to the main Keplarian orbit and is the same for all years, while the other component is due to something that varies over the data sets (Bertelsen, 2014).

The other possibility for the line differences is that B14 has a higher error in flux calibration than previously thought. If this is true, then S11 and B14 will match in line profile height, making only the wings differ. This difference in wings would indicate variability in S11 data which is especially prominent for mid and high J values. This is consistent with a single peaked constant component with a variable, broad component present in these line profiles (Bertelsen, 2014).

Bertelsen modeled HD 163296 in order to determine which scenario is more likely. A few results for different CO emission lines are shown in Figure 2.10. The model showed that the high J lines are in agreement with S11 for the FWHM, peak separation, and flux. B14 is also in agreement with the model for flux values of high J
lines. However, the modeled shape of the line flux versus the J curve does not match for either data set.

![Figure 2.10: Line profiles for CO v(1-0) P04, P14, and P37 modeled with ProDiMo radiation thermo-chemical disc code. The line wavelength, flux, FWHM, and peak separation are given on the plot (Bertelsen, 2014).](image)

The model predicts a much stronger double peak than seen in any of the NIRSPEC observations. This could be caused by either a low transmission at the center line of S11 or a non-Keplerian component present in both S11 and B14. Further work with the model shows that a non-Keplerian additional component in the low J line profiles from S11 and in all of the lines from B14 matches the data seen. Without the non-Keplerian addition, there are multiple inconsistencies between both sets of data. This non-Keplerian component is shown to be weaker in S11 and not present at all for the high J (Bertelsen, 2014).

A possible explanation for the variable, non-Keplerian component of the CO line profiles is a disk wind. A detailed discussion of disk winds and their effect on line profiles is given in the following section. Klaassen et al. (2013) found a low velocity, rotating molecular disk wind located within a few AU of HD 163296. This wind might have an outflow of CO which is causing the unusual shape for the line profiles. If this wind is not always constant, perhaps perturbed by the geometry of the disk, it would explain why the CO is less prominent in S11 and why the high J lines were not
affected here.

In order to test this theory, more measurements need to be taken. Due to the difficulties of observing NIR wavelengths with strong Telluric lines present, Neon fine structure lines might be an alternative way to detect the outflow of a disk wind. Ne II emission lines at 12.81\,\mu m can be seen out to 10-15 AU (Glassgold et al., 2007), and so it should be possible to determine if a molecular outflow causes variability in the lines close to the star, while remaining constant further out (Bertelsen, 2014).

2.4 Changes in Line Profiles

Many different authors have looked into possible causes for changes in CO and other line profiles originating near the inner rim of the accretion disk. This section will present several different possibilities for seeing unexpected shapes in CO emission.

2.4.1 Keplerian Rotating Disk

Pontoppidan et al. (2008) created a disk model of a Keplerian disk in order to compare observed line profiles to theory. This model assumes a flat disk, rotating with Kepler's laws, and a constant radial surface density. The disk is divided into rings with specific temperatures following a power law gradient.

\[ T = T_0 \left( \frac{R}{R_0} \right)^{-\alpha} \]

Where \( R \) is the radius, \( T \) is the gas temperature, and \( T_0 \) is the temperature at the inner radius \( R_0 \).

Bast et al. (2011) used this model to compare the CO line profiles of several different T Tauri and HAeBe stars to theory. They verified that CO profiles at low
inclination (4.3°) produce a single peaked, narrow line, while higher inclinations (70°) produced double peaked profiles. Data from TW Hya matched the low inclination model while VV Ser matched the high inclination.

However, no good fit was found for data from AS 2005 A. With an inclination of 22°, the model proved to be too narrow at the base, while too wide at the peak. Figure 2.11 shows the data and model of the three stars listed here. Several different inclinations were tried for AS 205 A, but none of them gave a good fit.

![Figure 2.11: Comparison between Keplerian rotation model and data from three separate stars (Bast et al., 2011).](image)

Further modeling work by Bast et al. (2011) shows that the line profiles can be modeled while using a non-standard temperature structure. Instead of assuming the power law temperature gradient, they looked for a temperature gradient which does give the correct profile. They were able to find good fits, but they involved including extended emission out to 30 AU. This extended emission results in a large velocity shift not seen in the data. It is concluded that a non-standard temperature is most likely not the cause of the differences between model and data (Bast et al., 2011).

### 2.4.2 Thermally Launched Winds

Thermally launched winds can be driven by FUV, EUV, or X-rays. This wind would result in the heating of the disk's atmosphere, causing the gas' thermal energy
to exceed the gravitational potential of the disk, producing a flow of material from the disk surface. The outflow from these winds will retain the Keplerian rotational velocity of the launch radius, resulting in a slight velocity shift between the star and the emission Bast et al. (2011).

If the winds are driven by EUV, the temperatures are thought to be so warm as to photodissociate any CO in the atmosphere. Further work needs to be done to confirm this, but for now Bast et al. (2011) excludes EUV thermally launched winds for this reason.

FUV and X-ray driven winds would result in cooler temperatures, but have a launching radii too far out in the disk. X-ray driven winds have a launch radius of 10-50 AU, while FUV is greater than 50 AU. However, soft X-ray winds have a much closer launch radius. Ercolano & Owen (2010) show that soft X-ray driven wind will result in a symmetric, broad-based, single-peaked line profiles with a small blueshift between 0-5 km/s.

2.4.3Magnetically Launched Winds

The protostar produces a powerful magnetic field which exerts a torque strong enough to launch winds from the disk. Edwards et al. (2006) studied these winds with He I λ10830. These winds resulted in very high velocity shifts in the line profiles which are not seen for any of the CO data.

2.4.4Funnel Flow

The magnetic field from the star is thought to be a link between the accretion disk and star via a funnel flow along the dipolar field lines (Hartmann et al., 1994). This funnel flow could be producing the unusual line profiles; however, Najita et al.
(2003) believes the temperatures inside the funnel to be 3000–6000 K higher than the measured temperatures of the CO-emitting gas and not capable of producing these kinds of line profiles. On the other hand, Bary et al. (2008) believes the funnel flow to be cooler where it connects with the disk. This region could be emitting the line profiles. If we assume the profiles are originating at the cooler region of the funnel flow, Najita et al. (2003) shows that a line emitted from this region will be double peaked with no velocity shift. While this is not the cause for Bast et al. (2011)’s CO line profiles, it could explain line profile discrepancies for other data sets.

2.4.5 Disk Wind and Keplerian Rotation

After these discrepancies surrounding CO came to light, Pontoppidan et al. (2011) created a new disk model with an additional disk wind component. This model includes an idealized wind structure constructed with a locus below the central star. The wind is briefly accelerated along a set of linear streamlines in order to determine how a line profile would change. An image of the disk and projected field lines are shown in Figure 2.12

This combined model readily produces single-peaked lines with broad wings for low inclinations. The wings are controlled by the Keplerian disk rotation while the peak is created from the non-Keplerian wind component originating within a few AU of the star. Figure 2.13 shows a comparison of the disk only line profile to the disk plus wind profile. If the wind originates from further out in the disk, a double-peaked profile is created. Also, the addition of a disk wind does not necessarily produce a blueshifted line.
Figure 2.12: Schematic of a Keplerian disk with disk winds originating below the disk. Several different wind directions are used in the model to determine effects at different radii (Pontoppidan et al., 2011).

Figure 2.13: Comparison of the modeled line profile with the disk only (blue) and with the disk plus a wind (red) (Pontoppidan et al., 2011).
The inclination of the disks effects how much of the wind is seen in the line profile. Figure 2.14 shows the resulting line profile for several different inclinations. As you can see, the more inclined disks produce a shallower, double peaked profile. HD 163296 has an inclination of $46 \pm 4^\circ$, as such we could expect a profile shape which is slightly double peaked and shallower if a disk wind is present.

![Figure 2.14: Comparison of the modeled line profile with various inclinations of the disk (Pontoppidan et al., 2011).](image)
Chapter 3

OH Emission

More information is needed to determine the cause and meaning of the variability seen in CO. However, we can take this opportunity to investigate previous observations that CO and OH are tightly linked in the inner regions of the disk (Brittain et al. 2016, Mandell et al. 2008). In this Chapter, we will review the excitation conditions of OH and what its presence near the inner rim indicates. OH spectra of HD 163296 is available for two different years, one of which coincides with the CO data from B14 (2012). We will compare the OH line profile and flux changes between 2009 and 2012 to those seen in CO.

3.1 Excitation Conditions

In order to understand the line profiles seen in stellar spectra, we must review the causes for emission from CO and OH molecules. Emission is the result of either electrons in the molecules gaining and losing energy through excitation, consequently absorbing and emitting photons, or the vibrational and rotational modes changing. Level populations are determined by spontaneous de-excitation, stimulated
emission/absorption, and stimulation through collisions. Spontaneous de-excitation occurs when a molecule randomly emits a photon, lowering its energy state. Stimulated emission/absorption is brought about by a photon of a specific wavelength interacting with an electron of a molecule, causing it to lower or raise its state respectively.

The states available to a molecule depend on the discrete energy levels decided by the total internal energy of a molecule. This energy depends on the sum of its rotational and vibrational energy.

\[
\frac{E_v}{\hbar c} = (v + 1/2)\omega - (v + 1/2)^2\chi\omega + B_v J(J + 1) - D_J J^2(J + 1)
\]

Where \(J\) is the total angular momentum of the molecule, \(B\) is the rotational constant, \(\omega\) is the angular velocity, and \(v\) is the vibrational quantum number. The population of various energy levels is given through the partition function.

\[
Z = \sum_j (2J + 1)e^{-E_j/kT}
\]

This shows that the population of a state depends on the energy and temperature of the molecule. In order for the excitations to be visible \(E_j/kT\) must approach unity. In regions of high temperature, (such as the inner rim of an accretion disk), the energy must also be high in order for transitions between states to occur.

For a diatomic molecule such as CO, the changes in \(J\) can be \(\Delta J = \pm 1\) and vibrational states can be \(\Delta v = \pm 1, \pm 2, \pm 3, \ldots\). This results in numerous possibilities of transitions between the three types of energy levels.

OH has a similar set of parameters, but with an added complexity of magnetic interaction between the electronic and vibrational levels. This results in hyperfine
splitting of the energy levels. This splitting causes each level to split into eight options instead of CO’s two options. A diagram of the energy levels of OH is shown in Figure 3.1.

Figure 3.1: Energy level diagram for the first two electronic states of OH. All transitions for the ground state with \( v = 0 \) and \( J = 5/2 \) are shown as well as all the transitions for the first excited state with \( v = 0 \) and \( J = 5/2 \) (Schleicher & Ahearn, 1982).

The nomenclature of Figure 3.1 uses X to represent the ground state, with the excited electronic states using A, B, C, etc. The rest of the terminology follows the
form.

\[ {^{2S+1} \Lambda^{+/-}}_{J,(g/u)} \]

Where \( S \) is the spin quantum number, \( \Lambda \) represents the orbital angular momentum given in values \( \Sigma, \Pi, \Delta, etc \), \( g/u \) is the parity, and +/- indicates the reflection symmetry. The rotational transitions between \( \Delta J = -1, 0, 1 \) are given the names R, Q, and P branch respectively (Schleicher & Ahearn, 1982). However, for both CO and OH \( \Delta J = 0 \) is forbidden and as such, no Q branch is possible. For this paper, we are only concerned with the ro-vibrational transitions within the ground state and P branch.

### 3.2 Photodissociation of Water

Water is present in warm accretion disks due to the combination of hydrogen and dust with oxygen. The three step process to create water is shown here

\[
H + H + \text{dust} \rightarrow H_2 + \text{dust}
\]

\[
\varnothing + H_2 \rightarrow OH + H
\]

\[
OH + H_2 \rightarrow H_2O + H
\]

Optically thick dust clouds shield hydrogen from UV radiation, blocking any UV photon absorption. However, dust does enhance molecular hydrogen formation by (1) providing a site for the hydrogen to meet, and (2) providing a sink for binding energy. This binding energy is then released and begins heating the dust grains and ejecting the \( H_2 \). As a result, molecular clouds are surrounded by shells of H I with the \( H_2 \) inside.
If molecular hydrogen is not present, water can also be created via OH and H

\[ \varnothing + \text{H} \rightarrow \text{OH} + \text{hv} \]

\[ \text{OH} + \text{H} \rightarrow \text{H}_2\text{O} + \text{hv} \]

Either way, the formation of gaseous water in the disk’s atmosphere strongly depends on the gas density (Fedele et al., 2014).

Photodissociation is an important way for molecules to change in a circumstellar disk. The photodissociation of water is how OH is created near the inner rim of the disk. When a photon of the appropriate frequency interacts with water, OH and H are created. This only occurs at temperatures well below 1000K. Above this temperature, photodissociation of OH occurs to create O and H.

\[ \text{H}_2\text{O} + \text{hv} \rightarrow \text{OH} + \text{H} \]

UV absorption of OH and H\textsubscript{2}O play an important role in the inner disk. In order to learn about where each of these molecules form and how they interact with material around them, Ádámkovics et al. (2014) created a model which accounted for the photodissociation of various molecules. This model showed that H\textsubscript{2}O is located below a warmer layer of OH, at small AU from the star. This coincides with theories that water is only found in deeper layers of the disk, shielded from photodissociation, where it is more difficult to observe (Fedele et al., 2014).
3.3 Uses of OH

CO is not the only molecular tracer that we can use to learn about the inner workings of accretion disks. OH line emissions have similar excitation energies as some of the CO emissions and can thus also be used to trace out the inner rim of the disk. Figure 3.2 shows where NIR interferometry is useful and which molecular lines are used for different regions of the disk.

![Figure 3.2: Schematic of an accretion disk with relevant wavelengths and molecules labeled (Dullemond & Monnier, 2010). This paper is only concerned with the inner regions of the disk.](image)

The presence of OH emission in disks is commonly associated with the flared disk geometry of Group 1 HAeBe stars. While HD 163296 is not alone in its abundance of OH (HD 85567), it appears to be rare for these types of stars (Fedele et al., 2014). Ro-vibrational lines of OH depend on the temperature of the surrounding material which, for accretion disks, depends on UV output. For the Group 1 stars, the inner rim and outer regions are both flared in such a way as to be thermally heated by the
star and receive abundant UV florescence. The Group 2 disks are shadowed by the inner rim and will hence receive less UV radiation. The presence of OH in a shadowed disk like HD 163296 indicates something else might be going on. A disk wind could be lifting the OH above the plane of the disk, causing the variable emission we see.

3.4 Instrumentation and Data Reduction

Archival data taken from the CRIRES (Cryogenic high-resolution InfraRed Echelle Spectrograph) instrument off of the European Space Observatory’s (ESO) Very Large Telescope (VLT) was used for one of the OH data sets and the CO data. CRIRES has a spectral range from 1 to 5.3 \( \mu m \) with a resolving power of up to \( \lambda/\Delta\lambda = 10^5 \) when using the 0.2 arcsec slit.

For the OH data, the telescope was nodded in an ABBA sequence in order to cancel the sky emission. This data had a 600 second integration time and was taken in May of 2012, less than three months after B14’s CO data in March of the same year. This means that we can assume that HD 163296 was in the same epoch and therefore, should not have a significant variance in the overall NIR brightness like we observe over the course of several years. The CO data was also taken in an ABBA sequence with a 60 second integration time.

The OH data from July of 2009 was obtained at the W. M. Keck Observatory with the NIRSPEC (Near Infrared Spectrometer) echelle spectrograph. This instrument operates in the spectral range 0.95-5.4 \( \mu m \), has a resolving power of \( \lambda/\Delta\lambda = 2.5^4 \) with a 0.432 arcsec slit, and was also nodded in an ABBA sequence.
<table>
<thead>
<tr>
<th>Instrument</th>
<th>CWLEN nm</th>
<th>Date</th>
<th>Integration Time</th>
</tr>
</thead>
<tbody>
<tr>
<td>CRIRES</td>
<td>2941</td>
<td>20120528</td>
<td>600s</td>
</tr>
<tr>
<td>CRIRES</td>
<td>4992</td>
<td>20120307</td>
<td>60s</td>
</tr>
<tr>
<td>CRIRES</td>
<td>4978</td>
<td>20120307</td>
<td>60s</td>
</tr>
</tbody>
</table>

### 3.4.1 OH Reduction

The CRIRES data was partially reduced using the ESO pipeline. Once the spectrum was extracted from the raw data, each frame was flat fielded, while hot pixels and scratches on the instrument were removed by hand. The same process was done for a standard star (eps Sgr) imaged the same night under similar telluric conditions. Once the spectra were clean, they were fitted to an atmospheric model to calibrate the wavelengths. This was done so that we could ratio out any telluric lines from the science data and observe the differences between what was expected and what was present for HD 163296. For the 2.9 $\mu$m range, a wide P4.5 OH line profile is found and shown in Figure 3.3.

The OH doublet from 2009 was already reduced when provided for this paper. The spectra of HD 163296 and the standard star along with the ratio is shown in Figure 3.4. The resolution for NIRSPEC is not as good as CRIRES and therefore the spectra do not fit quite as well.
Figure 3.3: HD 163296 spectrum around 2.9\(\mu m\) from 2012. Black represents the spectrum from HD 163296, while red is for the standard star. The residual found after ratioing the science data and the standard star is shown above the spectrum. OH line P4.5 can by seen at 3408 cm\(^{-1}\). A line was placed through the continuum to make it easier to see the line profile.

Figure 3.4: HD 163296 spectrum around 2.9\(\mu m\) from 2009. Black represents the spectrum from HD 163296, while red is for the standard star. The residual found after ratioing the science data and the standard star is shown above the spectrum. OH line P4.5 can by seen at 3408 cm\(^{-1}\). A line was placed through the continuum to make it easier to see the line profile.
3.4.2 CO Reduction

The data for CO line emission was also taken from ESO’s archive and is the same data used in B14. However, this data was reduced following the same process as the OH data for the sake of consistency between the two data sets. The standard star used for ratioing was Lam Sco. Figure 3.5 shows CO emission lines P26 and P27. For this paper, we will be comparing CO P26 to OH P4.5 due to their similar excitation energy.

![HD 163296 spectrum around 4.9µm from 2012](image)

Figure 3.5: HD 163296 spectrum around 4.9µm from 2012. Black represents the spectrum from HD 163296, while red is for the standard star. The residual found after ratioing the science data and the standard star is shown above the spectrum. CO lines P26 and P27 can be seen at 2032 cm\(^{-1}\) and 2027 cm\(^{-1}\).
Chapter 4

Analysis and Conclusions

In this Chapter, we will compare the line profiles of OH to determine how they changed over three years and then compare this to the variability seen in CO from Bertelsen (2014). This will indicate whether the variability in CO and OH could be due to the same mechanism. We will also review the possible reasons for variability in these molecules and discuss ways to verify these possibilities.

4.1 Results

For the following comparisons, the relevant line profiles have been extracted from the larger spectrum and corrected for Doppler shift. The two profiles are then plotted on top of each other with the residual shifted above.

4.1.1 OH Comparison

A comparison of OH doublet P4.5 is shown in Figure 4.1. The residual clearly shows the 2009 data to be double peaked, while the 2012 is much flatter. They both have similar FWZI values as shown in Table 4.1, but the FWHM differs greatly. This
difference is due to the OH doublet being clearly defined for 2009 and not 2012. The presence of a disk wind in 2012 could be the cause of the flattening of OH in 2012. As the model indicated in Section 2.4.5, the inclination of HD 163296 would cause line profiles to shallow out and develop slightly double peaked profiles.

Table 4.1: Comparison of the FWHM and FWZI from manual measurements of OH data taken in 2012 and 2009. All numbers are given in km/s.

<table>
<thead>
<tr>
<th></th>
<th>CRIRES 2012</th>
<th>NIRSPEC 2009</th>
</tr>
</thead>
<tbody>
<tr>
<td>FWHM</td>
<td>70±15</td>
<td>54±15</td>
</tr>
<tr>
<td>FWZI</td>
<td>120±15</td>
<td>120±15</td>
</tr>
</tbody>
</table>

Figure 4.1: Comparison of OH from 2012 and 2009. Blue represents the 2012 data from CRIRES, while Black represents 2009 data from NIRSPEC. The residual is shifted above the profiles and a line is placed along zero flux to help clarify the differences.
4.1.2 CO Comparison

Figure 4.2 compares the OH P4.5 doublet from 2012 to the CO P26 line. Here, the OH and CO profiles are scaled based on their equivalent widths in order to make comparison easier. This Figure shows the CO data to be peakier than the flat OH data. However, the widths of CO and the OH doublet at zero intensity are very similar. This is shown in Table 4.2. (Note: The values given are for the entire OH P4.5 doublet).

This comparison shows the OH doublet and single CO line emission to have very similar shapes. In order to confirm whether this is realistic or not, an idealized Gaussian version of a CO line profile and an OH doublet were each convolved with a wide Gaussian emission created from disk rotation. The convolution of CO shows that the narrow, idealized line emission fully emulates the disk rotation once convolved. The OH doublet results in a slightly broader convolution. This excess in the wings will be mostly lost in the noise of real data. This proves that the comparison of OH to CO in Figure 4.2 is realistic.

<table>
<thead>
<tr>
<th></th>
<th>CO P26</th>
<th>OH P4.5</th>
</tr>
</thead>
<tbody>
<tr>
<td>2012</td>
<td>60±15 (km/s)</td>
<td>54±15 (km/s)</td>
</tr>
<tr>
<td>FWZI</td>
<td>120±15 (km/s)</td>
<td>120±15 (km/s)</td>
</tr>
<tr>
<td>Eq Width</td>
<td>0.033 (cm(^{-1}))</td>
<td>0.0069 (cm(^{-1}))</td>
</tr>
</tbody>
</table>
Figure 4.2: Ratio of OH doublet P4.5 and CO P26 from 2012. Red represents the OH data while black indicates CO. The CO emission was scaled down with the equivalent widths to match OH. The residual is shifted above the profiles and a line is placed along zero flux to help clarify the differences.

Figure 4.3: The top Figure shows the convolution (red) of a narrow gaussian representing CO (blue) and a broad emission resulting from disk rotation (black). The convolution and broad emission appear directly on top of one another. The bottom Figure shows the convolution of the same broad emission with the idealized OH doublet. This results in slightly broader wings. The difference between each convolution and broad emission is shown in black above the profiles.
4.1.3 Other Herbig Ae Be OH/CO ratios

Brittain et al. (2016) is working on an extensive study of ro-vibrational emission of OH and CO from Herbig Ae/Be stars. This study shows the relationship OH has with other aspects of the protoplanetary disk by comparing the OH luminosity to UV, PAH, stellar temperature, and CO luminosities. They found that the OH doublet and high J CO lines have a relatively constant ratio. However, more work needs to be done to determine whether there is a distinct difference in the Group I and Group II ratios. Figure 4.4 shows the relationship of OH luminosity to CO luminosity for 21 different HAeBes from Brittain et al. (2016) as well as HD 163296.

Table 4.3: Star ID numbers from Brittain et al. (2016).

<table>
<thead>
<tr>
<th>No.</th>
<th>Star</th>
<th>No.</th>
<th>Star</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>HD 76534</td>
<td>12</td>
<td>HD 45677</td>
</tr>
<tr>
<td>2</td>
<td>HD 259431</td>
<td>13</td>
<td>HD 85567</td>
</tr>
<tr>
<td>3</td>
<td>HD 250550</td>
<td>14</td>
<td>Elias 3-1</td>
</tr>
<tr>
<td>4</td>
<td>HD 100546</td>
<td>15</td>
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<tr>
<td>5</td>
<td>HD 31293</td>
<td>16</td>
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<td>6</td>
<td>HD 179218</td>
<td>17</td>
<td>HD 34282</td>
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<tr>
<td>7</td>
<td>MWC 765</td>
<td>18</td>
<td>HD 190073</td>
</tr>
<tr>
<td>8</td>
<td>HD 36112</td>
<td>19</td>
<td>UX Ori</td>
</tr>
<tr>
<td>9</td>
<td>HD 135344b</td>
<td>20</td>
<td>BF Ori</td>
</tr>
<tr>
<td>10</td>
<td>LkHα 224</td>
<td>21</td>
<td>CO Ori</td>
</tr>
<tr>
<td>11</td>
<td>HD 142527</td>
<td>22</td>
<td>HD 163296</td>
</tr>
</tbody>
</table>
Magnitude data in the M and L bands from Malfait et al. (1998) was used to determine the luminosity of OH and CO of HD 163296. This data gave a flux density of the star which allowed the calculation of the line flux for each emission line. We found the line flux of OH P4.5(+1) to be $1.18 \times 10^{-14}$ ergs/s/cm² while the line of flux of CO P26 is $8.95 \times 10^{-14}$ ergs/s/cm². This gives a ratio of 7.5, consistent with Brittain et al. (2016)’s results.
4.2 Conclusions and Future Work

The comparison of OH between 2009 and 2012 proves to be inconclusive. There certainly appears to be a distinct difference in line shape, but the signal to noise is simply not high enough to know for sure. The variation appears to be consistent with the type of variability seen in CO (Bertelsen, 2014), but further measurements need to be taken in order to confirm this trend.

However, the comparison of OH and CO from 2012 shows apparent similarity. The similar FWHM and FWZI shows an important relationship which needs to be tested by further measurements. If the OH continues to match CO throughout several years of data, it would indicate that the same driving force is causing variability in each molecule.

Overall, more data is needed to determine the cause of the variability seen in HD 163296. We need to first determine whether the CO variability is indeed caused by disk winds. A model needs to be created with HD 163296’s specific inclination and other parameters. Once this is done, the effects of a disk wind at various points in the disk needs to monitored and compared to the archived data. Also, the non-Keplerian component assumed to be responsible needs to be confirmed by taking more data in the M band at multiple time intervals. This will help us to understand how the wind might be changing over the decades and allow us to discern the natural state of CO line emission in the disk.

Once the cause of CO variability is determined, it will be easier to tell whether or not the same cause is possible for OH variability. Again, more data needs to be taken in the L band at various time intervals in order to discover whether the variance is periodic or random. If the OH data proves to change with similar intervals as CO, we can assume that the same driving force is responsible for both molecules.
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