VLT/X-shooter observations of the massive young stellar object B275 in M17

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Abstract

We have analyzed the VLT/X-shooter spectrum of the candidate massive young stellar object (mYSO) B275 in the star forming region M17. Observing the formation of massive stars is difficult, because the process takes place on a short timescale and is obscured by many magnitudes of optical extinction. With sensitive observations of B275, we have firmly classified the spectrum of a mYSO for the first time, providing an accurate temperature. Subsequently, the luminosity of B275 was determined by de-reddening the SED and B275 was placed in the HR-diagram. A large discrepancy between optical classification and luminosity was found. Even though the presence of the first and second CO overtone emission suggests a hot and dense (accretion)disk near the stellar surface, no evidence for ongoing accretion was seen. In this thesis, the current evolutionary status of the peculiar system of B275 will be discussed.
Acknowledgements

I would like to thank my supervisor, Lex Kaper, for his clever insights and everlasting motivation during the past two years, Lucas Ellerbroek, for his dedicated assistance during the research of this thesis, and the ESO staff, for obtaining the X-shooter spectrum.
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1 Introduction

Stars are the building blocks of the observable Universe and how they form is one of the central questions in modern astrophysics. They power the luminosity of galaxies and the first generations of stars have been responsible for the reionization of the Universe. The heavy elements we find on Earth and in the human body are a direct consequence of the cycle of stellar evolution, as these elements are formed deep inside stars through nuclear fusion and are returned to the interstellar medium when the stars end their live as planetary nebulae or supernovae, providing material for the formation of the next generation of stars. The birth of stars is also inevitably linked with the formation of planetary systems.

Mass is the most important factor in the properties and evolution of a star. It determines the size, luminosity and even the lifetime is constrained by the amount of mass the star contains. Due to their higher luminosity, massive stars deplete their central hydrogen much faster compared to their lower mass counterparts. The ultimate fate of a star is also determined by its mass, because only the most massive stars (M > 8 M⊙) will explode as a supernova and eventually become a neutron star or a black hole, whereas lower mass stars shed their outer envelopes through a planetary nebula and end up as a white dwarf.

Even though small in number, massive stars play a dominant role in the physical, chemical and morphological evolution of galaxies. They interact with the surrounding interstellar medium (ISM) through dense stellar winds, massive outflows, expanding H II regions and the eventual collapse and supernova explosion.

![Figure 1: Left: the IRDC G11.11-0.11 seen against the bright mid-IR emission of the galaxy. Right: The H II region NGC6618 in M17.](image)

The first observable stage of star formation is an infrared dark cloud (IRDC, figure 10.2.6). IRDCs are cold, dense molecular cloud cores, which are seen in silhouette against emission of the galaxy in the mid-infrared. Observations of the interior of these darks clouds reveal that they can contain density clumps and the most massive density clumps are believed to be the precursors of massive star formation. When the temperature inside the cloud reaches ~ 100 K,
complex molecules evaporate of dust grains and emit a broad spectrum in the sub-millimeter area. This stage is often referred to as the 'hot core' phase. As the temperature rises further, the protostar will ionize its immediate surroundings until the object becomes visible at radio and infrared wavelengths, due to the heating of dust and/or the (infrared) hydrogen recombination lines. This stage is observable as an ultra-compact H II region (UCH II). The ionized region around the young star or cluster expands and disrupts the parent molecular cloud until it becomes a classical H II region (see figure 10.2.6). At this stage, the stellar content of the star forming region becomes detectable at optical and near-infrared wavelengths.

One of the biggest H II regions in our Galaxy is the Messier 17 (M17) complex. M17, also known as the Omega Nebula or Swan Nebula, is located in the constellation of Sagitarrius. It was discovered by Philippe Loys de Chéseaux in 1745 and catalogued as M17 in 1764 by Charles Messier. Being a Messier object, it is one of the brightest nebulas on the night sky and therefore one of the best laboratories in the Galaxy to study star formation in a rich, massive cluster.

The gas in the H II region is excited by a cluster of OB stars embedded in the adjacent M17 south west (M17SW) giant molecular cloud with a total luminosity of $L \sim 5.6 \times 10^6 L_\odot$ (Harper 1976 [1]). The total mass of the system, including the adjacent M17SW giant molecular cloud (receding at a velocity of 20 km s$^{-1}$), exceeds $3 \times 10^4 M_\odot$ (Lada 1976 [2]). The open cluster NGC6618 is deeply embedded in the M17SW cloud. It contains about 35 massive stars and is the driving source of the ionization, resulting in the nebular spectrum.

Observations of M17 have led to a better understanding of the physical conditions in giant molecular clouds and photodissociation regions (the boundary of the ionized H II region with the neutral molecular gas). For a long time, the content of the NGC6618 cluster and other massive star forming regions was unknown, but recent significant improvement in near-infrared instrumentation has advanced our understanding of massive stars and their early evolution.
2 Massive star formation

Our knowledge of the formation of low to intermediate mass stars is understood in great detail compared to their massive counterparts. For stellar masses up to $\sim 8 \, \text{M}_\odot$, the formation is explained by the nonhomologous collapse of a (slowly rotating) dense part of the parent molecular cloud. The central parts will rapidly form a core, while the surrounding dust-gas envelope continues to fall inward. Eventually, the object will become optically thick and this will halt the collapse temporarily. At a temperature of $T \geq 2000 \, \text{K}$, molecular hydrogen dissociates and a second phase of collapse begins. The collapse starts in the centre of the cloud, where this temperature is first reached and propagates outwards. Therefore, it is called an 'inside out collapse'. The core will continue to accrete material, while contracting on a Kelvin-Helmholtz timescale (see section 3), which will increase the temperature. At $T \sim 10^7 \, \text{K}$, hydrogen will ignite and the star will reach the zero age main sequence (ZAMS). The system loses angular momentum through a circumstellar disk and/or outflows.

One of the main problems in our understanding of the formation of massive stars is that the radiation pressure of stars with a mass greater than $\sim 8 \, \text{M}_\odot$ is high enough to reverse the flow of material and stop further accretion onto the stellar surface (Larson & Starrfield 1971 [3]). These theories were proposed decades ago and involved spherical geometry. Furthermore, stars with spectral type as early as B ionize hydrogen and thus create an H\text{ii} region around them. This ionized region has a temperature and pressure two times that of the surrounding molecular gas, making it even more complicated for an accretion flow to reach the stellar surface (Keto & Wood 2006 [4]). In addition to these theoretical complications, the observational study of the most massive stars is hindered by the following points:

1) All massive stars, with the exception of a few regions such as the Orion star forming region, are located at large distances ($\geq 1 \, \text{kpc}$). Sensitivity and angular resolution of observations are not sufficient to determine the geometry and flow direction of the accreting material.

2) As the birth and early evolution of massive stars takes place in the dense inner parts of molecular clouds, observations of these evolutionary stages are limited by the large amount of dust extinction. Only long wavelength photons (sub-mm, radio) are able to penetrate the parental cloud.

3) Due to the short timescales on which massive stars evolve, the chance to observe a massive star in formation is small. For example, Iben (1965) [5] calculated that the pre-main sequence (PMS) phase for a 15 $\text{M}_\odot$ star is $6 \times 10^4 \, \text{yr}$, while a 1 $\text{M}_\odot$ star stays a pre-main sequence object for $5 \times 10^7 \, \text{yr}$. As these calculations assumed a fixed mass, the luminosity of the PMS star determines the timescale for gravitational contraction. Palla & Stahler (1992) [7] calculated that for protostars with an accretion rate of $10^{-5} \, \text{M}_\odot \, \text{yr}^{-1}$, stars more massive than 8 $\text{M}_\odot$ already start to burn hydrogen while accreting and therefore have no PMS phase.
These points are only partially compensated by the fact that massive stars have high luminosities and can therefore be studied at larger distances. Because of the lack of observational evidence, theoretical models have remained controversial. Recent observations and theoretical models hint towards the fact that massive stars do form in a similar way to low-mass stars, i.e. through disk accretion, despite the effects of the radiation pressure on the circumstellar accretion flow. This should eventually lead to a single or stellar central system. In fact, the majority of massive stars seem to form in binary or multiple systems (Preibisch et al. 2001 [6]). In recent literature, three main models are proposed for massive star formation:

1) **Monolithic collapse:** A single protostellar core accretes material from its gravitational sphere of influence, which ultimately leads to the accumulation of mass at the center of the core. The final mass will depend on how the material has been distributed amongst possible members of the forming system.

2) **Competitive accretion:** In this case, the gravitational influence is more extended, as multiple cluster members pull the molecular cloud gas inwards. The protostars also move relative to the gas, which increases the reservoir of available material. The only gas which is truly destined for a particular star is within its envelope and circumstellar disk; the remaining gas will fall down to the deepest point in the gravitational potential well.

3) **Stellar mergers and collision:** Larson (1971) [3] showed that the radiation pressure on dust particles is too great for stars more massive than 8 M\(_\odot\) to continue accretion. Also, the apparent clustering of massive stars is dense, such that a sufficiently large gas reservoir for the formation of a cluster of massive stars would not be available. According to Larson [3], the only way for high mass objects to form is through merging of multiple members in dense stellar clusters (Baumgardt & Klessen 2010 [9]).

The primary difference between monolithic collapse and competitive accretion is that during monolithic collapse, most of the material is gathered before actual star formation occurs. In competitive accretion, the material is acquired during stellar formation through an active mechanism, which keeps matter flowing from the molecular cloud towards the center of the cluster (the deepest point in the gravitational potential well). To estimate growth through stellar collisions, we calculate the cross section \(\sigma\) of two stars with mass \(M\), passing each other at a minimum distance \(R_{\text{min}}\). The cross section is enhanced due to gravitational focusing:

\[
\sigma = \pi R_{\text{min}}^2 \left(1 + \frac{2GM}{R_{\text{min}}v^2}\right)
\]

Assuming for the moment a velocity dispersion for stellar clusters of 10 km s\(^{-1}\) (Fabian et al. 1975 [10]), the second term in de bracketes dominates. With the stellar collision rate we can estimate the timescale between stellar collisions:
\[ \tau_{\text{coll}} = \frac{1}{n_{\text{star}} \sigma_{\text{grav}} v} \]  

\( n_{\text{star}} \) represents the stellar number density. With typical representative parameters for a dense stellar cluster we obtain:

\[ \tau_{\text{coll}} = 7 \times 10^7 \left( \frac{n_{\text{star}}}{10^6 \text{pc}^{-3}} \right)^{-1} \left( \frac{M_*}{10 M_\odot} \right)^{-1} \left( \frac{R_{\text{min}}}{1 R_\odot} \right)^{-1} \left( \frac{v}{10 \text{km s}^{-1}} \right)^{-1} \text{yr} \]  

The velocity dispersion \( v \) is one-dimensional and for extremely young clusters this dispersion can be larger than 5 km s\(^{-1}\) (Mengel et al. 2002 [8]). The most massive stars known remain on the main sequence for \( \sim 3 \) Myr. Therefore, the most massive stars will have ended their life before they encounter another massive star with the observed properties of a typical cluster. This threshold timescale could decrease to the order of \( \tau_{\text{coll}} \sim 10^6 \) yr, in case one would include the possibility of nonequal mass encounters and the increase of the collisional cross section due to circumstellar disks. Simulations by Baumgardt & Klessen (2010) [9] show that the theory of stellar mergers would in fact require clusters to be too compact compared to observed massive clusters. Their conclusion is that stellar collisions only play a minor role in the formation of massive stars.

### 2.1 Theory of massive star formation

The basic sequence of massive star formation can be summarized as: (1) The contraction and (2) the collapse of a giant molecular cloud into a hot core, which (3) subsequently accretes material from its surroundings and eventually becomes visible as it (4) disrupts the birth cloud.

#### 2.1.1 Giant molecular clouds

The ISM of galaxies contains gas that is inhomogeneous; the physical conditions of the gas span a wide range, from hot (10\(^6\) K) and tenuous X-ray emitting plasma to cold (10 - 100 K) and dense molecular gas. This molecular gas is of vital importance for star formation. \( \text{H}_2 \) is the main component of the gas and is accompanied by a rich chemistry of complex molecules.

Giant molecular clouds (GMCs) are composed primarily of \( \text{H}_2 \), but direct observation of the \( \text{H}_2 \) gas is difficult. Radiative transitions of \( \text{H}_2 \) are weak, owing to the molecule’s lack of a permanent dipole moment. In addition, the lowest lying rotational energy levels of \( \text{H}_2 \) are widely spaced and therefore rarely excited in GMCs with typical temperatures of 10 - 15 K (McKee 1999 [13]). The next most abundant molecule is \( \text{CO} \) (\( \sim 2 \times 10^{-5} \) as abundant as \( \text{H}_2 \)) and is often used as a tracer of the distribution of \( \text{H}_2 \).

It is surprising that such molecules can exist in the ISM, particularly because of the destructive effects of ultraviolet radiation coming from massive stars. Atomic hydrogen shields most interstellar gas from energetic photons with energies 13.6 \( \leq h\nu \leq 100 \) eV, but photons with energy 5 \( \leq h\nu \leq 13.6 \) eV
Figure 2: The distribution of CO in the Milky Way. This figure is taken from Dame et al (2000) [11]. Nearby regions, such as the Orion complex (right), span a larger range in galactic latitude. Most of the molecular gas is located within \( \sim 4 \) kpc of the galactic centre. The sun orbits at a distance of \( \sim 8.3 \) kpc around the galactic centre, which is located in the direction of Sagittarius A* (\( l = 0^\circ \)).

penetrate deep into molecular structures. Molecules are able to form if their absorption lines are optically thick, or when the molecules are shielded by interstellar dust. A value of \( A_V \geq 0.08 \) (Savage 1977 [12]) towards a cloud surface is typically sufficient for a molecular cloud to gain a substantial column density of molecules (i.e. \( N(H_1 + H_2) \sim 5 \times 10^{20} \) cm\(^{-2}\)).

Giant molecular clouds are the largest structures in the Galaxy. Their physical properties have been discussed by McKee (1999) [13]: sizes range from \( \sim 20 - 100 \) pc, containing masses of \( 10^4 - 10^6 \) M\(_\odot\). The typical temperature is 10 - 15 K and the average local density \( \rho_H \) is on the order \( 4 \times 10^3 - 1.2 \times 10^4 \) cm\(^{-3}\). On the other hand, the volume averaged density is estimated at \( \sim 50 - 100 \) cm\(^{-3}\), which indicates that the clouds are highly 'clumped', probably due to turbulent motions inside the cloud, which compress local pockets of gas.

2.1.2 Contraction and collapse

The distribution of CO in the Milky Way is plotted in figure 2. Molecular clouds are concentrated in the centre and along the spiral arms of our Milky Way. The spiral arms are thought to originate from density waves propagating through the plane of the galaxy (Lin & Shu 1964 [14]). In turn, these waves compress local gas and dust environments. The interaction with a shock front coming from a group of massive stars (originating from either an expanding H II region, or a supernova explosion) would be another way of compressing a cloud. This mechanism is often referred to as ‘triggered star formation’ (Elmegreen & Lada 1977 [15]).

For a cloud to collapse, gravity has to deal with several factors: (1) the gas pressure (2) magnetic forces (3) internal turbulence (4) rotation. A simple approximation (where the magnetic forces, rotation and internal turbulence of the cloud are neglected) to the collapse of a cloud is given by the Jeans mass:

\[
M_J = 1.1M_\odot \left( \frac{T_{\text{gas}}}{100 \text{K}} \right)^{\frac{3}{2}} \left( \frac{\rho}{10^{-19} \text{ g cm}^{-3}} \right)^{-\frac{1}{2}}
\]  

(4)
The Jeans mass is defined as the smallest mass which is stable against gravitational collapse. If \( M > M_{\text{J}} \), the cloud is unstable against its own gravity and collapses. In equation 4, typical values for a molecular cloud are given. Supersonic turbulence inside the cloud as a repulsive force will die out when it is not constantly being injected with energy. When the cloud is assumed optically thin, i.e. radiation created through for example recombination can escape the cloud without interacting, the gas collapses on a free-fall timescale:

\[
t_{\text{ff}} = 2.1 \cdot 10^5 \text{yr} \left( \frac{\rho}{10^{-19} \text{g cm}^{-3}} \right)^{-\frac{1}{2}}
\]

(5)

This shows that the densest parts of the cloud will collapse faster. The cloud will therefore fragment in smaller structures. These dense parts heat up when they become optically thick and the formation of a star(cluster) could commence.

2.1.3 Accretion and disruption

Stars with mass \( M > 8 M_\odot \) will have enough radiation pressure to stop accretion, when spherical symmetry is assumed (Larson 1971 [51]). This follows from equating the outward radiation pressure to the inward gravitational attraction:

\[
\frac{F_{\text{rad}}}{F_{\text{grav}}} = \left( \frac{\kappa L}{4\pi r^2 c} \right) / \left( \frac{GM}{r^2} \right)
\]

(6)

\( \kappa \) is the opacity of the circumstellar material, \( L \) the luminosity, \( r \) the radius (which cancels out) and \( c \) the speed of light. Furthermore, stars of spectral type B and earlier are hot enough to ionize hydrogen and create an \( \text{H}^\text{II} \) region. The ionized gas will have a temperature and pressure about a factor of two higher than the molecular environment. The increased pressure of the ionized gas will hinder the accretion process and the remaining protostellar material will expand back into the ISM. To illustrate the difficulties of massive star formation even more, we note that for early O stars a very high accretion rate is needed (\( \dot{M} > 10^{-3} M_\odot \text{yr}^{-1} \)), as the star would otherwise have evolved of the main sequence before the accretion process has finished (Hosokawa 2010 [16]).

The following conditions determine if the (proto)stellar object is able to maintain an accretion flow:

1) The effective opacity \( \kappa_{\text{eff}} \) of the accretion flow could be variable, or even reduced. When the object is embedded, the emerging radiation field is largely absorbed at UV and optical wavelengths by the dust particles and re-emitted at (far)IR wavelengths. The gas is more transparent at these wavelengths. Density inhomogeneities in the circumstellar material could reduce \( \kappa_{\text{eff}} \). The radiation can escape at moments when the accreting material is optically thin, or when the accretion flow seizes for a certain time interval.
2) The impacting flux can be reduced by the non-spherical geometry of the accretion flow. If the circumstellar material is put into an accretion disc, the radiation can escape in the polar directions (see figure 3). A circumstellar disk is a natural consequence of angular momentum conservation during the collapse of a cloud and is therefore expected to determine the geometry of the accretion flow near the star. The high density in the disk will shield the particles from the radiation field as they move closer to the star. Eventually the particles pass through the disk and face the UV and optical radiation from the stellar object. Dust is destroyed at \( T \sim 1500 \text{ K} \), or the particles must be contained in larger aggregates, so that the opacity at these distances is dominated by gas rather than dust. Observationally, it will be difficult to detect disks around massive stars, because the high UV-flux will photo-evaporate disks on timescales of \( 10^5 \text{ yr} \) (Hollenbach et al. 2000 [17]).

3) The gravitational acceleration can be enhanced with respect to the radiative acceleration. In this case, massive stars are formed in a cluster of low luminosity sources. The density in this cluster of low-mass objects peaks in the molecular cloud if \( \rho_{\text{objects}} \gg \rho_{\text{material}} \). The formation of massive stars could be increased by the coagulation of multiple low mass stars in these dense clusters. If this scenario is the only way to produce massive stars, the observation of an isolated massive star would be very rare.

During the accretion process, the central (proto)star will emit an increasing amount of ionizing photons as it becomes hotter and more luminous. In addition to this, a powerful stellar wind starts to interact with the surrounding material, destroying the accretion disk. A possible geometry of a mYSO is shown in figure 3. Close to the star, where the escape velocity exceeds the sound and/or flow speed, material will stay bound. Material in the polar directions will be cleared through a combination of a stellar wind and radiation.

During the pre-main sequence evolution, there are phases where the circumstellar material has a non-spherically distribution (Liseau 2004 [19]). Often hour-glass shaped mass outflows are observed and/or wide bi-polar or narrow jets are formed. This process is believed to play a major role in the angular momentum budget of the star-disk-jet system. Stellar jets are powered by gravitational energy, which is liberated by the accretion process of the disk material. Evidence for this picture comes from the strong correlation of outflow signatures (P Cygni profiles, forbidden line transitions) and accretion diagnostics (UV and IR emission excesses, or inverse P Cygni profiles). The decline in outflow activity as a function of stellar age and disk frequency shows us that a disk and a jet are well connected to the earliest phases of star formation. Jets are powered from a wide range in radii from the circumstellar disk through photo-evaporation (‘disk wind models’), or through the interaction of the inner region of the disk with the stellar photosphere (‘X-wind models’), as discussed in Shu et al. (1995) [20]. These outflows are magnetically collimated into a jet. Subsequently, the jets interact with the infalling material and collide with
the surrounding cloud, where they cause a shock and inject turbulence into the ISM. The mechanical energy of the outflow after impact is converted into heat and can excite certain molecules, most prominently H$_2$.

2.1.4 Final products of massive star formation

The final products of high mass star formation are OB clusters, OB associations, or possibly field OB stars.

1) OB clusters, such as the Orion nuclear cluster NGC36-4, are relatively dense star clusters with 1 - 100 massive O- or B- type stars and a mass of $10^3$ - $10^5$ M$_\odot$. The central part becomes more concentrated as more massive stars sink to the core due to mass segregation.

2) OB associations are star forming regions, where massive O- and B- type stars are spread over large distances in the parent molecular cloud. One example is the Sco OB 2 cluster. OB associations are different to OB clusters, which are typically created in a small volume. Distances range from 1 to 10 parsec between different candidate massive stars. It is believed
that stellar formation progresses gradually through these regions, due to interaction of the radiation field of young stellar objects with molecular clouds (Blaauw 1964 [22]).

3) Although it is still under debate, massive OB stars could form in isolation. These are often called ‘field OB stars’, as they are found outside clusters or associations. These stars should not be mistaken with runaway OB-stars, which are ejected from their birth location with velocities > 40 km s\(^{-1}\). Some of the field stars can not be linked to a nearby cluster or association, which is a textbook example of our lacking knowledge in the formation of the most massive stars (De Wit et al. 2004 [21]).

2.2 The initial mass function: an upper limit to star formation?

The initial mass function (IMF) is an empirically determined relationship that describes the initial mass distribution of a population of stars.

\[
\frac{dN}{d\log M} \propto M^{-x}
\]

The number of stars with masses in the range \(M + dM\) within a specified volume is proportional to \(M^{-x}\) (\(x\) being a dimensionless exponent). Salpeter (1955) [23] first proposed a power law index of -1.35 and the average distribution of stellar masses seems to follow this index very well. Table 1 shows the amount of stars found within a certain mass interval, using different logarithmic slopes. These numbers are normalized in such a way that there is exactly one star present in the highest mass interval. It illustrates how rare massive stars are.

<table>
<thead>
<tr>
<th>Mass range</th>
<th>(x = 1)</th>
<th>(x = 1.35)</th>
<th>(x = 1.7)</th>
</tr>
</thead>
<tbody>
<tr>
<td>0.5 - 1 M(_\odot)</td>
<td>128</td>
<td>700</td>
<td>3822</td>
</tr>
<tr>
<td>1 - 2 M(_\odot)</td>
<td>64</td>
<td>275</td>
<td>1176</td>
</tr>
<tr>
<td>2 - 4 M(_\odot)</td>
<td>32</td>
<td>108</td>
<td>362</td>
</tr>
<tr>
<td>4 - 8 M(_\odot)</td>
<td>16</td>
<td>42</td>
<td>111</td>
</tr>
<tr>
<td>8 - 16 M(_\odot)</td>
<td>8</td>
<td>16.6</td>
<td>34.3</td>
</tr>
<tr>
<td>16 - 32 M(_\odot)</td>
<td>4</td>
<td>6.5</td>
<td>10.6</td>
</tr>
<tr>
<td>32 - 64 M(_\odot)</td>
<td>2</td>
<td>2.6</td>
<td>3.3</td>
</tr>
<tr>
<td>64 - 128 M(_\odot)</td>
<td>1</td>
<td>1</td>
<td>1</td>
</tr>
</tbody>
</table>

Table 1: Number of stars calculated in specified mass intervals using different logarithmic slopes \(x\).

Determining the IMF for massive stars is complicated, as they tend to lose mass quickly through outflows and stellar winds. Furthermore, massive stars are often born in the center of clusters. The mass segregation of the cluster forces the massive stars to these crowded parts and even when they can be detected, they
more than occasionally turn out to be binaries or multiple systems. One key question is if the IMF is determined at the earliest stages of cluster formation (i.e. the fragmentation of the cloud prior to collapse of dense parts) or that it is shaped in later processes (competitive accretion or after the enrichment from nearby stars).

In the past, numerous studies were aimed at the question whether there is an upper limit to the formation of stars. Through the IMF, the maximum mass of a star should depend on the mass of the parental molecular cloud. For example, a $10^3 \, M_\odot$ cloud can only form an $8 \, M_\odot$ star, whereas a $10^5 \, M_\odot$ cloud is able to create a $50 \, M_\odot$ star. Most massive stars found to date do not exceed $130 > M_\odot > 140$, which led to a general belief of a firm upper mass limit of $\sim 150 \, M_\odot$ (Figer 2005 [24]). However, several sources of the R136 star cluster seem to exceed the accepted mass limit (Crowther et al. 2010 [25]). As these stars drive powerful winds, it is clear that the opacity of the ionized gas is high enough for the radiative acceleration to exceed the surface gravity. It is possible that the upper limit for star formation is set by the mass loss rate of a (proto)star becoming greater than the accretion.
3 The pre-main sequence evolution of massive stars

After the collapse of the parent molecular cloud, the protostellar object will accrete material from its surroundings. When the accretion phase ends and the envelope disperses, the object will become visible and becomes a pre-main sequence (PMS) star. The PMS star acquires its luminosity through gravitational contraction. This contraction will occur on a Kelvin-Helmholtz timescale:

$$\tau_{KH} = \frac{E_{\text{int}}}{L} \approx \frac{E_{\text{gr}}}{2L} \approx \frac{GM^2}{2RL} \approx 1.5 \times 10^7 \left( \frac{M}{M_\odot} \right)^2 \left( \frac{R}{R_\odot} \right)^{-1} \left( \frac{L}{L_\odot} \right)^{-1} \text{yr} \quad (8)$$

During this contraction, the central temperature rises and when hydrogen burning starts, the star has reached the main sequence. The first PMS tracks have been available from Iben (1965) [5]. In this theory, PMS stars of all masses derive their luminosity solely by the gravitational contraction of the object. In the early phases of this contraction, the protostar could have a radius nearly two orders of magnitudes larger than the final ZAMS value. It is now known that these protostars are not able to reach these large radii during the gathering of mass from the molecular cloud (Larson (1969) [51]). Palla & Stahler (1990 [26]) predicted that stars would rather evolve off a well-defined 'birthline' by combining the classical theory of PMS evolution with numerical calculation of the mass-radius relation of accreting proto-stars. The birthline would be the locus in the HR-diagram where the accretion stops and the star becomes optically visible.

In figure 4, the birthline is shown corresponding to a constant accretion rate of $\dot{M} = 10^{-5} M_\odot \text{yr}^{-1}$ and $\dot{M} = 10^{-4} M_\odot \text{yr}^{-1}$. The increase in accretion rate will shift the birthline vertically in the HR-diagram if the stars accrete the same amount of material. A higher accretion rate will then lengthen the PMS tracks. A star of 1 $M_\odot$ will finish accretion well before it reaches the main sequence. However, stars which continue to accrete material for a long time will start burning hydrogen before they reach hydrostatic and radiative equilibrium and therefore will not have a proper PMS phase. For an accretion rate of $10^{-5} M_\odot \text{yr}^{-1}$, the birthline will coincide with the ZAMS at $M \sim 8 M_\odot$. The distribution of the PMS Herbig Ae/Be stars correlates well with the theoretically derived birthline for $M = 10^{-5} M_\odot \text{yr}^{-1}$.

A YSO is invariably associated with large amounts of gas and dust and the formation of the most massive stars may well proceed through the accretion of material through a circumstellar disk. Evidence for this scenario is accumulating by the detection of disks around massive young stellar objects (mYSOs). To date, the largest accretion disk observed has a diameter of $\sim 20,000$ AU (Chini et al. 2004 [27]), which is about 250 times the size of our Solar System. The mass of the central object is estimated at $\sim 20 M_\odot$. The most recent model on the formation of massive stars through disk accretion is published in Hosokawa et al. (2010) [16].
Figure 4: Pre-main sequence tracks for non-accreting protostars from Iben (1965), shown in solid lines. The upper heavy solid line is the birthline corresponding with an accretion rate of \( M = 10^{-4} \, M_\odot \, \text{yr}^{-1} \) (Palla & Stahler 1990), while the lower heavy solid line is for \( M = 10^{-5} \, M_\odot \, \text{yr}^{-1} \). In this plot, the main sequence is not drawn. The squares are Herbig Ae/Be stars (Finkezeller and Mundt (1984)). The filled symbols are outflow sources from Levreault (1988).
4 The star forming region of M17

The M17 complex is one of the biggest H II regions on the sky. Scientific research of M17 has long focussed on the gas and dust in the luminous H II region or the molecular cloud around it, but the stellar content powering the H II region has remained a mystery for a long time. Due to the large line of sight extinction, decent UV and VIS spectroscopic observations were very difficult to achieve. Therefore, the first massive stars were identified in M17 based on JHK photometry. Follow-up K-band spectroscopy can then be used for spectral classification. For massive stars, these estimates are not as precise as spectral classification using UV or optical spectroscopy, where a hot photosphere produces more lines. In this section we will discuss observations, parameters and the stellar content of the M17 region.

4.1 Structure

Figure 5: The M17 field (in galactic coordinates) taken with GLIMPSE (red: 8.0 μm, orange: 5.8 μm, green: 4.5 μm, blue: 3.6 μm). The boundaries of the large-scale structures are overlaid. The NGC6618 cluster and the methanol maser source G15.03-068 (used for the trigonomical parallax distance determination [33]) are located in the M17 South region at the boundary with the H II region. This figure is adapted from Povich et al. (2009) [29].

The image shown in figure 5 reveals the structure of the M17 complex. The H
The II region divides the M17 molecular cloud in two halves (M17 North and M17 South). Early CO measurements (Elmegreen 1976 [15]) showed an extensive molecular cloud structure associated with M17 towards the southwest (SW) and showed that UV photons could penetrate deep into the molecular cloud. This was attributed to an internal ‘clumpy’ structure and it was suggested that the entire complex is undergoing star formation, beginning with the stars in the NGC6618 cluster at the north east tip of the structure adjacent to the H II region. The boundary of the molecular cloud/H II region interface is oriented nearly edge-on, which makes M17 a favorable object to study the effects of UV radiation in molecular clouds. The northeast side of M17 interacts with the M17EB region that contains several OB type stars. However, this cluster is separate from NGC6618.

With the improved sensitivity of telescopes and instruments, some spectra of the less reddened objects could be obtained by Hanson et al. (1997) [28]. In this work, 6 O-type stars and 7 mYSOs were identified and over 100 B-type stars were uncovered in Povich 2009 [29].

Large line of sight extinction does not necessarily mean that the object or cluster is very young. M17 was classified as an extremely young star forming region because of the high number of massive stars observed in the region and the spectroscopic and/or photometric signatures of circumstellar material contained inside the cluster. The most massive stars in M17 predominantly agree with a 1 Myr isochrone (Hanson et al. (1997) [28]).

<table>
<thead>
<tr>
<th>M17</th>
<th></th>
</tr>
</thead>
<tbody>
<tr>
<td>Distance</td>
<td>1.3-2.1 kpc</td>
</tr>
<tr>
<td>RA2000</td>
<td>18:20:26</td>
</tr>
<tr>
<td>DEC2000</td>
<td>-16:10:36</td>
</tr>
<tr>
<td>App. mag. (V)</td>
<td>6.0</td>
</tr>
<tr>
<td>App. dim.</td>
<td>11 arcmin</td>
</tr>
<tr>
<td>Luminosity</td>
<td>$5 \times 10^6 , L_\odot$</td>
</tr>
<tr>
<td>Mass</td>
<td>$&gt; 3 \times 10^4 , M_\odot$</td>
</tr>
</tbody>
</table>

Table 2: M17 parameters

### 4.2 Extinction and distance

The first measurements of the extinction towards M17 showed an east-to-west gradient, starting at $A_V = 1$ mag at the eastern H II region and peaking at $A_V = 7$ mag at the central dust filament in front of the cluster. The most recent and detailed measurements so far yielded an $A_V$ of 2-6 mag at the eastern, optically bright part of the nebula, which is consistent with Galactic models at a distance of 2 kpc as presented in Bahcall & Soneira 1980 [30]. Hanson et al. (1997) [28] reported an average value $A_V = 8$ mag. Clearly, the extinction towards M17 is variable on small spatial scales. It is composed of the Galactic extinction with $R_V = 3.1$ for the total-to-selective extinction, plus an intrinsic value inside M17, with a possible higher R-value (Hoffmeister & Chini 2008 [32]). $R_V$ is defined as follows:

$$R_V = A_V / E(B-V)$$

With $E(B-V) = A_B - A_V$. It is thought that the value of $R_V$ is determined by
the local dust size distribution (Cardelli, Clayton & Mathis (1989) [42] (CCM)).

While M17 contains one of the closest H ii regions in the Galaxy, the distance towards the cluster is still a matter of debate. UBVRI photometry of Chini et al. (1980) [31] showed that M17 is located at 2.2 kpc based on dereddening of several sources and estimation of spectral types. Hanson et al. (1997) [28] classified the optical spectra of several early type objects and corrected them for extinction using the K-band magnitudes rather than in the optical. As the extinction in the K-band is one tenth of that in the V-band (on the magnitude scale), the error in de-reddening will be significantly lower. After correction, the authors compared the magnitudes with zero-age main sequence (ZAMS) parameters to determine a distance of $1.3^{+0.4}_{-0.3}$ kpc. Hoffmeister & Chini (2008) [32] used three different methods to derive a distance of 2.1 kpc. The 'method of variable extinction' is used to investigate deviations from the general galactic extinction law, characterized by the value of $R_V = 3.1$. The results are visualized in figure 6.

The slope of the fit corresponds to the correct $R_V$ value inside M17, while

Figure 6: Distance modulus $V - M_V$ versus the colour excess $E_{B-V}$. The sample consist of 53 OB stars in M17 with $A_V > 2$. Filled circles are stars earlier than B3, open circles represent stars later than B3. The foreground extinction with $A_V = 2$ mag and $R_V = 3.1$ is shown as a dashed line in the upper left, where stars with $A_V < 2$ are rejected in order to exclude foreground stars (dotted line). This figure is taken from Hoffmeister et al. (2008).
the extrapolation towards $E_{B-V} = 0$ is due to the foreground extinction at $A_V = 2$ and the galactic $R_V = 3.1$ value. The derived distance modulus of 11.6 mag reveals a distance of $2.1 \pm 0.2$ kpc towards the cluster. Secondly, the calculated reddening law at $R_V = 3.9$ and the foreground extinction of $A_V$ were used to list the distance moduli of 45 cluster members. Again, $2.1 \pm 0.2$ kpc was concluded. Finally, the luminosity of the whole cluster (from the spectral energy distribution) and the contribution from the OB-stars were scaled such that $L_{SED} = L_{OB}$. This yielded a minimum distance of $d = 2.1$ kpc.

Recently, Xu et al (2011) [33] measured the distance towards M17 using the radio parallax of G15.03-0.68 in M17, resulting in a distance of $1.98^{+0.14}_{-0.12}$ kpc which we will adopt.

4.3 Stellar content

Figure 7 is taken from Hanson et al. (1997) [28] and shows the infrared color-color diagram J-H versus H-K of M17. The solid line in the lower left represents the position of the main sequence at a distance of 1300 pc, ranging from spectral types O3 V to M III. The dashed lines are the reddening vectors for an O3 V and M III star, respectively. The vectors show the track of ‘normally’ reddened stars without an infrared excess or giant atmosphere. Only sources with K-band magnitudes $K \geq 10$ mag are plotted. A lot of information can be drawn by just looking at the plot:

1) Stars with normally reddened photospheric colors are clearly seen. These stars would fall between the dashed lines.

2) The amount of infrared excess sources to the right of the dashed line.

3) The position in the diagram provides an estimate of the reddening of the observed sources.

The naming convention of the sources comes from Bumgardner (1992) [34]. Hanson et al. (1997) [28] made a survey of all the known (at that time) OB-stars in the southwestern star forming region of M17 and in figure 8 the locations of the OB and the candidate mYSOs are plotted on top of a DSS image of M17. Note that some objects are circular as well as squared. These candidate mYSOs were classified using optical spectra as an OB star by Hanson et al. (1997) [28]. The object of interest in this project is the infrared excess source B275.
Figure 7: The near-infrared color-color diagram of M17. Marked in a square box is our object of interest M17B275. The solid line in the lower left represents the main sequence for unreddened stars. Using the standard extinction law given by Koornneef (1983) [37], ‘normally’ reddened stars should fall between the two dashed extinction lines. This figure is taken from Hanson et al. (1997) [28].
Figure 8: A 9.5’ x 7’ field of view of M17 from the DSS server. North is up, East is left. Plotted are the locations of the OB stars (circles) and candidate mYSOs (squares) in M17 taken from Hanson et al. (1997) [28] in M17SW. B275 is plotted in red. The stars are all located in the western part of the nebula where the dense local dust environment blocks the light at shorter wavelengths causing these stars to be heavily obscured.

4.3.1 B275

Hanson et al. (1997) [28] detected no photospheric features in B275 in the blue, besides hydrogen Balmer lines and the 4430 Å diffuse interstellar band. Two different scenarios were proposed to explain the lack of lines: B275 may undergo veiling, either due to a continuum source or in the individual lines. Another possibility was that because of the young nature of the object, B275 could have a rotational speed close to the breakup speed, which would make the weak stellar features hard to detect. Therefore, the nature of this source remained uncertain. Further spectral analysis by Hanson et al. (1997) [28] has revealed that B275 shows significant CO first overtone emission, as well as heavy Brγ emission. The CO emission is likely to originate from a rotating disk, as the CO (2-0) first overtone displays a blue shoulder (see section 10.2.6). Thought to be a B0 V star according to the extinction corrected flux (Hanson et al. (1997) [28]) and listed to be a B2 V (Simbad), there was no consensus on the actual evolutionary status of B275. We present a new UVB to NIR spectrum, taken during the X-shooter first science verification run, which reveals the true nature of B275.
5 Observations & data reduction

For this project, spectra obtained with the ESO Very Large Telescope and the new medium resolution spectrograph X-Shooter (e.g. Vernet et al. 2011 [35]) were analyzed. The observations were carried out on the night of August 11 2009 with the VLT UT2 8.2 meter telescope during the first science verification run of X-shooter (P.I. R. Chini, Ruhr-Universität Bochum). X-shooter has a unique spectral coverage from the UV (300 nm) to the NIR (2500 nm) in a single exposure. All science images were taken and read out within one hour and split into multiple exposures.

The VLT is located in the Atacama desert of Chili on mountaintop Paranal (see table 3), where disturbing factors such as light pollution and turbulence are minimal. It is very important to have optimal observing conditions in order to have good quality data. The object should be as high as possible above the horizon to minimize airmass and the position of the moon should be taken into account when scheduling observing time.

The illumination of the moon during the observation is significant (76%), but the distance between the moon and the object was large enough to carry out the observations (figure 9). At UT=03:00, the airmass was 1.08. A two dimensional Gaussian function was fitted to the optical acquisition image (see figure 10) to determine the quality of the images. The full-width-at-half-maximum (FWHM) of the Gaussian is a good measure for the atmospheric seeing at the time of observation. Several non-saturated stars were taken and the FWHMs in the x and y direction were calculated. The observations have a seeing of 0.58" in the V-band. The average seeing in the period 1999-2006 at Paranal was 0.83", therefore, the observing conditions were excellent. Our object in these conditions was smaller than the slit width and thus we were seeing limited (see table 4 for the resulting spectral resolution).

In table 4 the calculated parameters of our observations are shown. As our object is heavily reddened, the integration time in the UVB must be much longer than in the NIR. The total integration time for NIR is split into 12 different frames. Each frame contains 20 detector integration times (NDIT) of 11 seconds (DIT) to prevent saturation of the detector. This means that the detector is read out after every 11 seconds, as further collection of photons in the detector will not be possible, resulting in loss of signal.

All observations were acquired by nodding the star on the slit using a ABBA...
sequence. This allows for background subtraction using the sky spectrum on position B when the target is in position A (and vice versa). For a full description of nodding sequences, we refer to the official X-shooter pipeline manual (Modigliani et al. 2011 [36]). Due to the high line-of-sight-extinction in M17, it is difficult to obtain a decent signal in the UVB (300 - 550 nm) spectrum. Observations are binned 1x2, to create a pixel with two times more signal, but with only single readout noise. This will increase the signal-to-noise ratio, though still oversampling (see table 4). The right choice of slit width is also important when dealing with regions of high extinction and, as a result, a large slit width was chosen in the UVB observations to gather more light into the spectrograph at the cost of resolution. In table 4 the resolutions as shown in the official X-shooter manual ($R_{\text{theo}}$) are compared with the values of the ones obtained during the science verification run ($R_{\text{obs}}$). The observed resolution was calcu-
lated using the xsh_wavecal recipe contained in the X-shooter pipeline through the Gaussian fitting of a large number of lines. The FWHM of these Gaussians yield the resolution of the spectrum, which is defined as:

\[ R = \frac{\lambda}{\Delta \lambda} \]  

where \( \Delta \lambda \) represents the minimum wavelength scale where the spectrograph can still distinguish between two different spectral features. \( R \) is the dimensionless resolving power. The pixel sampling corresponds to the number of pixels covering one resolution element. To represent the resolution element in velocity we can use:
Table 4: Instrument setting during M17B275 observations. The sampling is given in pix/FWHM.

<table>
<thead>
<tr>
<th>Arm</th>
<th>$t_{int}$ (s)</th>
<th>$\lambda$ (nm)</th>
<th>slit (&quot;)</th>
<th>$R_{theo}$</th>
<th>$R_{obs}$</th>
<th>Samp.</th>
<th>S/N</th>
</tr>
</thead>
<tbody>
<tr>
<td>UVB</td>
<td>$4 \times 685$</td>
<td>300 - 550</td>
<td>1.6</td>
<td>3300</td>
<td>2696</td>
<td>4.9</td>
<td>$\sim 74$</td>
</tr>
<tr>
<td>VIS</td>
<td>$8 \times 285$</td>
<td>550 - 1050</td>
<td>0.9</td>
<td>8800</td>
<td>7658</td>
<td>6.9</td>
<td>$\sim 85$</td>
</tr>
<tr>
<td>NIR</td>
<td>$12 \times 20 \times 11$</td>
<td>1050 - 2500</td>
<td>0.9</td>
<td>5600</td>
<td>5681</td>
<td>4.2</td>
<td>$\sim 60$</td>
</tr>
</tbody>
</table>

\[ \frac{\lambda}{\Delta \lambda} = \frac{c}{\Delta v} \]  \hspace{1cm} (11)

$c$ is the speed of light and $v$ our resolution on the velocity scale. This yields $\Delta v_{UVB} = 90 \text{ km s}^{-1}$, $\Delta v_{VIS} = 34 \text{ km s}^{-1}$ and $\Delta v_{NIR} = 54 \text{ km s}^{-1}$. The binning in the dispersion direction of the UVB arm reduces the sampling by a factor of 2, but this does not affect the resolution element. The VIS (550 - 1050 nm; unbinned) has a wavelength step of 0.02 nm and the NIR (1050 - 2500 nm; unbinned) has a wavelength step of 0.06 nm. Finally, the last column of table 4 shows the signal-to-noise ratio of the observations and was calculated using the MIDAS STATIS/IMAGE command and obtained through:

\[ S/N = \frac{\text{mean}}{\text{SD}} \]  \hspace{1cm} (12)

The signal-to-noise level is equal to the mean value of the signal divided by the standard deviation. Several measurements were taken and the average represents the value of the signal-to-noise ratio. Note that this ratio is variable across the spectrum. For the flux calibration of the UVB and VIS spectra the standard white dwarf star EG274 (spectral type DA2) was observed. The flux standard EG274 was observed with the maximum slitwidth of 5" to avoid slit losses. Then the response curve of the X-shooter was obtained by dividing the observed spectrum by a table containing the true fluxes from the ESO website. A flux calibrated spectra of B275 is then obtained by dividing the reduced science frames with the response curve. We also corrected for slit loss, integration time and binning in the UVB. In the NIR the telluric standard HD180699 (spectral type B8 V) was used for the flux calibration and telluric line removal. With the use of the IDL routine Spextool, the stellar hydrogen lines were removed from the spectrum. The telluric continuum was then scaled to a blackbody with a temperature of an A0 V star which was fitted to magnitudes from literature. In this way the response curve of the instrument was obtained in the NIR. After correction for slit loss, the flux calibrated spectrum was calculated by dividing the science frames with the response curve. Telluric line removal was done with the Xtellcorr routine within the Spextool package. The raw science frames were bias-subtracted, flatfielded, corrected for cosmic ray hits and wavelength calibrated using the X-shooter pipeline version 0.9.4. Analyses of the reduced images were performed with IRAF, MIDAS and IDL.
6  The X-shooter spectrum of B275
6.1 Spectral features

Thanks to the wavelength coverage and sensitivity of X-shooter the spectrum of B275 has been obtained in unprecedented detail. While previous observations only revealed Balmer lines and the DIB feature at 4430 Å (Hanson et al. (1997) [28]), numerous spectral features have now been resolved. This allows for accurate spectral classification (section 7), analysis of the interstellar spectrum (section 8) and the study of circumstellar material (section 10). The spectrum includes information on:

1) **Photospheric spectrum:** The photospheric spectrum shows deep and broad Balmer, Paschen and Brackett hydrogen absorption lines and He I. Metal absorption lines include Mg II, C II and Si II.

2) **Interstellar spectrum:** We detected 16 diffuse interstellar bands (DIBs). The interstellar atomic Ca II K and Na I D doublet transitions are also prominent in the line of sight towards B275.

3) **Emission line spectrum:** The strongest hydrogen lines are seen in emission. Hα and Hβ show double peaked profiles, although these are likely to be an artifact of the nebular emission subtraction. The Ca II triplet and the O I 8446 Å line show a clear double peaked profile. The higher transitions of each series are also partially or completely filled in with emission. In the near-infrared, the first (2.3 µm) and second (1.5 µm) CO overtone emission features are detected.

4) **Forbidden lines:** The [O I] 6300 and 6363 Å are detected as well as some weak forbidden emission lines of iron. These lines show no blue shifted component that would be indicative for an outflow and are detected along the whole slit. They are therefore likely to be of nebular origin.

Order merging problems appear in the VIS spectrum in the 5500 - 7500 Å range. The Balmer-jump at ∼ 3650 Å and the Paschen jump at ∼ 8200 Å are also seen, although they have been normalized and removed from the spectrum. The flux calibrated spectrum (discussed in section 9.2) corroborates the infrared excess, beginning at ∼ 1 µm and shows both the Balmer and Paschen jump.
7 Spectral classification

In this section, the observations of B275 and the classification of the spectrum according to the Morgan-Keenan (MK) spectral classification system (Morgan & Keenan (1973) [38]) are discussed. The MK system is an iterative classification system, based on the observed spectrum. A set of standard stars defines the system and classification is done by visual comparison, taking into account all spectral features.

At the time the MK system was defined, the photographic emulsions were sensitive only to violet-blue light. Therefore, the classification was based on photographic spectra in the violet-blue part of the spectrum (3800-4600 Å). It contains a high concentration of temperature dependent lines in hot OB-stars, but the classification of cool stellar atmospheres is better done at longer wavelengths.

7.1 Spectral type

Figure 11: Main sequence stars, ranging from spectral type B2 to B8. The spectra in this section are obtained at the Dark Sky Observatory, acquired using the Gray/Miller Spectrograph, mounted on the 0.8M telescope of the Appalachian State University. The resolution element is 1.8 Å, close to the resolution of our observations of B275 ∼ 1.3 Å. Note the presence of the interstellar Ca ii K line and the DIB at 4430 Å.

The strength of the Balmer absorption lines are useful in assigning a spectral type in first iteration. As Hα and Hβ show strong emission from the hot circumstellar gas (as will be thoroughly discussed in section 10), the higher Balmer lines are used, because these are least affected by emission due to their origin in deeper photospheric layers.

The absence of doubly ionized He shows that the star is not hot enough to be of O-type. Furthermore, the star is excluded to be of early B-type, as the He i 4009 and C ii 4267 lines are very weak. These lines should be prominent down to a spectral type of B3 V.
The \( \text{He} \, \text{I} \, 4471 / \text{Mg} \, \text{II} \, 4481 \) ratio is one of the most useful spectral indicators for mid-to-late type B-stars (Gray & Corbally [39] (2009)). The neutral helium lines decay throughout the B spectral series and disappear around spectral type A0, whereas the Mg high excitation line (\( \sim 9 \) eV) grows in this regime. The dominant broadening mechanism of the helium line is pressure broadening, as in the Balmer series lines. This effect is caused by the interaction of the neutral atoms with electrons and ions in the star. This would explain the increasing strength of the Balmer sequence (reaching its maximum at \( \sim \) A0), as the amount of ionization drops towards the later type B-stars. Note that in the ratio \( \text{He I} \, 4471 / \text{Mg II} \, 4481 \), the equivalent width (EW) is taken and not the intensity of the lines, because the EW is (almost) independent of rotation and spectral resolution.

If the strength of the \( \text{Si II} \, 4128 / \text{He I} \, 4144 \) lines is also considered, we classify the photospheric spectra as B6. The luminosity class will be discussed in section 7.2. However, the spectrum does show peculiarities when classified as a B6 star. The strength of the Balmer lines differ from the standard stars in the spectral range B2-B8. Deep absorption cores are seen, which would indicate a spectral type later than B6. In addition, the He lines seem weak for a B6 atmosphere. Nevertheless, by judging the other spectral features, the star has to be of mid-B type.

The high signal-to-noise spectrum gave us the opportunity to make the first accurate spectral classification of B275. Hanson et al. (1997) [28] suggested that M17B275 was suffering photospheric veiling due to accretion on the YSO, which would explain the non-detection of photospheric features. This would produce both optical and UV continuum emission, that would fill in the stellar photospheric absorption lines. In case of optical veiling, the Balmer lines would be even deeper and they would therefore suggest a later type star. However, the \( \text{He I} \, 4471 / \text{Mg II} \, 4481 \) ratio should be independent of veiling as the continuum radiation is wavelength independent and, consequently, no significant veiling in the blue-violet is seen.

### 7.2 Luminosity class

Even though M17 is a young cluster (\( \sim 1 \) Myr old (Hanson et al. (1997) [28])), it is useful to derive a luminosity class using the broadening of the spectral lines. To exclude helium anomalies, the first thing to check is if the helium absorption lines correspond with the strength of the Balmer lines. There are several useful ratios of spectral lines available to constrain the luminosity class. The \( \text{Si III} \, 4552 / \text{He I} \, 4387 \) lines continues to be useful until spectral type B5, but weakens with decreasing luminosity (see figure 12). In our case, the \( \text{Si III} \) line at 4552 Å is not visible, which leaves the Balmer lines as the best indicator for the luminosity class in late B-type stars.

Because of the decreasing pressure in line forming regions of giant stars, the width of the spectral lines is smaller. The Balmer line profiles are affected by the drop in pressure.
Figure 12: Luminosity effects at B5. The line profiles become much "sharper" with increasing luminosity (and thus: radius) and lower log g.

Figure 13: From top to bottom: H\(\gamma\), H\(\delta\) and H\(\epsilon\). The spectrum of B275 (solid line) is overplotted with a B5 main sequence model (long dashed line) and a B5 giant model (short dashed line).

The effect on the Balmer line shapes is subtle in the luminosity range III-V. The B275 Balmer lines are deeper, even though they are likely to be filled in with circumstellar emission (as is shown in H\(\alpha\) and H\(\beta\)). When looking at the line wings, the H\(\gamma\) and H\(\epsilon\) follow the giant profile slightly better than the main sequence profile. H\(\delta\) does not fit very well, especially in the core, and the wings follow the main sequence profile better. The FWHM of the Balmer lines are \(\sim 1800 \text{ km s}^{-1}\) and the \(v \sin i = 114\pm11 \text{ km s}^{-1}\) (see section 7.3). Therefore,
the effect of rotation should be insignificant at these widths. The Inglis-Teller formula, where the number of detected Balmer lines is related to the electron density, is not used due to the fact that higher (and weaker) Balmer lines are lost in the noise at < 3700 Å.

The fact that B275 shows characteristics of a giant star is surprising, as the star is embedded in a young star forming cluster and these stars are not expected to have evolved of the main sequence. However, it may be that B275 is still contracting towards the main sequence on a Kelvin-Helmholtz timescale. The giant features reveal that the pressure is still relatively small and has not yet reached its main sequence value.

Using FASTWIND (Puls 2005 [45]), multiple models with different temperatures T were generated. These models corresponded to B5, B6, and B7 atmospheres (respectively 15250, 13750, 12750 K), with log g 4.0 and 3.5 (main sequence or giant star atmospheric pressure). The B7 III model gave the best fit profiles and confirmed the observed giant features in the spectrum of B275.

Figure 14: The giant sequence of B-type stars.

Figure 15: Left: Line profiles of the H\(\epsilon\) line in B275 (solid lines) fitted with different main sequence models, generated with FASTWIND (Puls 2005 [45]). Right: same, but for giant models.
Figure 16: Left: Line profiles of the HeI 4471 line in B275 (solid lines) fitted with different main sequence models, generated with FASTWIND (Puls 2005 [45]). Right: same, but for giant models.

7.3 Rotational velocity

The projected rotational velocity $v \sin i$ (with $i$ the inclination of the rotation axis with respect to the line of sight) is an important parameter during the evolution of a star. For example, the internal chemical composition can change significantly due to mixture induced by the rotation. Due to the resulting Doppler broadening of the spectrum lines, the value of $v \sin i$ can be measured. Alternatively, one can compare the profiles to those of a set of standard stars with known spectral type and rotational velocity; this gives a $v \sin i$ determination within an accuracy of $\pm 10\%$. The set of standard stars is defined in Slettebak (1975) [40]. The spectrum lines used for the comparison of $v \sin i$ values depend on spectral type and should meet the following conditions:

1) The lines are sufficiently strong and thus not lost in the stellar continuum in case of rapid stellar rotation.

2) Broadened by rotation only.

3) Unblended.

For the late O- and B-stars, only the hydrogen and helium lines are strong enough. Unfortunately, these lines are heavily Stark-broadened. We chose the He I 4471 Å line for the spectral range O9-B8. The line is also blended with the forbidden [He i] 4470 Å line. This forbidden transition is associated with a mixing of upper states in transitions of He I, which leads to a transition that is normally disallowed by the selection rules for electric dipole transitions (Beauchamp 1997 [41]). Together with Stark broadening of the He I 4471 Å line, this was taken into account when calculating the theoretical line profile for different amount of stellar rotation. With the use of a theoretically calculated line profile, the $v \sin i$ values of a set of standard stars were determined using the FWHM of the He I 4471 Å line.
Table 5: Interpolation of the $v \sin i$ parameter through the two standard stars 30 Sex and $\alpha$ Sco located in the constellations Sextans and Scorpius, respectively.

Table 5 lists the FWHM values of the standard stars and of B275. These standard stars were taken, as they were the closest match in spectral type compared with B275. The approximate $v \sin i$ value of $114 \pm 11$ km s$^{-1}$ was calculated through the interpolation of these values. This value was also used in our FASTWIND models and the fact that the calculated line profiles reproduce the width of the spectral lines well confirmed the estimated $v \sin i$.
8 The diffuse interstellar bands & ISM column density

In spectra of objects behind a significant column density of interstellar material, many absorption features (∼ 200) are seen, whose carrier(s) are still unknown. These absorption lines are called diffuse interstellar bands (DIBs), named after the likely origin of these mysterious stellar features, the diffuse ISM. DIBs were not initially recognized as interstellar, but after strong proof that they did not respond to the Doppler velocity shift and the fact that DIB strength roughly correlated with the reddening E(B-V) towards a background object (Herbig et al. (1975) [47]), it has been evident that these features are interstellar.

Although the carriers of (almost all) the DIBs are unknown, they have proven to correlate with certain species, such as atomic hydrogen (see figure 19). In this section, the detection of DIBs in the spectrum of B275 will be discussed. In particular, two of the most extensively observed DIBs (5780Å and 5797Å) allow to estimate the column density of neutral hydrogen in our line of sight.

Because of their dependence on color excess, it was thought that DIBs were formed on the surface of interstellar grains. These particles cause the amount of color excess in our line-of-sight. The interstellar reddening law, expressed in equation 9, deviates from its mean Galactic value \(R_V = 3.1\) in for example star forming regions. Hoffmeister & Chini (2008) [32] concluded a value of \(R_V = 3.9\) for M17, suggestive of a dust distribution which differs from the Galactic average. Herbig (1993) [48] showed that the DIB strengths did not correlate with an increase of the \(R_V\) value, which provided one of the first reasons to doubt if DIBs are created on grain surfaces. Some species of free polyatomic molecules, such as fullereen (C\(_{60}\)) and polyaromatic hydrocarbons (PAHs), are now favored candidates.

DIBs are extremely broad (0.8 - 30 Å) compared to atomic (inter)stellar line profiles. Velocity- or pressure effects do not affect the DIBs in any way and, therefore, these widths are probably due to the unresolved rotational structure of the carriers.

The equivalent width of the unresolved DIB features in the spectrum of B275 are listed in table 6. The error in measuring the equivalent width has several factors (continuum placement, signal-to-noise, blending with (inter)stellar features) and is at the 10% level. Hanson et al. (1997) [28] measured the strengths of the 4428 and 4502 Å features of several YSO sources inside M17 and the values we derive for B275 correspond very well to Hanson et al. (1997) [28] (see table 7).

Table: Detected DIBs.

<table>
<thead>
<tr>
<th>DIB</th>
<th>Strength (Å)</th>
</tr>
</thead>
<tbody>
<tr>
<td>4428</td>
<td>2.35</td>
</tr>
<tr>
<td>4502</td>
<td>0.26</td>
</tr>
<tr>
<td>5780</td>
<td>0.63</td>
</tr>
<tr>
<td>5797</td>
<td>0.14</td>
</tr>
<tr>
<td>6203</td>
<td>0.40</td>
</tr>
<tr>
<td>6613</td>
<td>0.21</td>
</tr>
<tr>
<td>6993</td>
<td>0.12</td>
</tr>
<tr>
<td>7432</td>
<td>0.77</td>
</tr>
<tr>
<td>7562</td>
<td>0.05</td>
</tr>
<tr>
<td>8026</td>
<td>0.22</td>
</tr>
</tbody>
</table>

Even though the visual extinction in the sample of Hanson et al. (1997) [28] varies significantly, i.e. \(A_V = 3-10\) mag, the DIB features do not follow this trend. If the DIB features are mostly tracing the foreground dust, which
Table 7: Comparison between YSO inside M17 of the equivalent width $W_{\lambda}$ and FWHM of the 4428 and 4502 DIB feature. All measurements, except for B275, are taken from Hanson et al. (1997) [28].

would reveal that the carrier is of interstellar origin, they are not affected by the local environment of M17. Furthermore, if the DIB features already saturate at a few $A_V$, additional absorbing molecules are unable to increase the EW. Circumstellar material (see section 10) could also affect DIB strengths, which has been observed in T-Tauri and other pre-main sequence objects (Herbig 1993 [48]). It has been assumed that the presence of hot dust and viscosity effects in circumstellar disks could reduce the abundance of DIB carriers, i.e. with respect to the measured $A_V$.

Cox et al. (2005) [49] found that the diffuse interstellar bands at 5780, 5797 and 6613 Å correlate well with the reddening in the galactic line of sight. B275 can be included in the figures plotted by Cox et al. (2005) [49] after measuring the strengths of the DIBs in the X-shooter spectrum. The measured values are given in equivalent width (EW), expressed in mA.

Figure 17: From left to right: The DIBs at 5780, 5797 and 6613 Å.

The value for the reddening, predicted by the 5780, 5797 and 6613 Å DIB strengths, yield an average $E(B-V) = 1.1$ (depending on the type of sightline; see figure 18). With our choice of $R_V$, ranging from 3.1 - 3.9, the visual extinction would be estimated at $3.4 \leq A_V \leq 4.3$.

As noted by Herbig (1975) [47], the DIB at 5780 Å follows the column density of atomic hydrogen particles $N(H_1)$ rather than the total column density, which is composed of both atomic hydrogen $N(H_1)$ and molecular hydrogen $N(H_2)$. In the absence of dust, which applies for the diffuse ISM, hydrogen is largely
composed of $N(H\text{ i})$. This is because $H_2$ is known to be produced on the surface of dust grains.

Figure 18: 5780 (top left panel), 5797 (top right panel) and 6614 Å DIB strength versus color excess, adapted from Cox et al. (2005) [49]. The sample is taken from several sources, as indicated in the plots. The linear fits exclude data of heavily reddened objects. Whether our line of sight passes through an edge- or core dominated cloud determines the source as a $\sigma$ or $\zeta$-type object, respectively. B275 (solid line) intersects the dashed lines at the measured DIB EW, which corresponds to an average value of $E(B-V) = 1.1$.

The relationship between $\log EW(5780 \text{ Å})$ and $\log N(H\text{ i})$ is shown in figure 19. The correlation with neutral atoms does not imply that the DIB carrier should be uncharged. Unfortunately, measurements on elements with similar degree of ionization (e.g. Ca, Ti, Mg) are very difficult, because of the low abundance of these elements in the ISM.

In figure 19, a $\sigma$ or $\zeta$-type are pointed out. These sources show significantly less scatter around the linear regression compared to the reddening $E(B-V)$, shown in figure 18. This implies that the DIB at 5780 Å follows the gas tracer $H\text{ i}$ much better than the dust tracer $E(B-V)$. With a combined sample from Cox et al. (2005) and Herbig (1993) [48], a correlation is found between neutral hydrogen and the DIB 5780 Å strength:

$$\log EW(DIB5780) = -16.4 + 0.87 \log(N_{H\text{ i}}) \quad (13)$$
Figure 19: The DIB strength at 5780 Å versus $N(H\ i)$ column density. This figure is taken from Cox et al. [49] (2005)

According to this relation, the neutral hydrogen column density towards B275 can be estimated at $\log N_{\text{HI}} = 21.94 \pm 0.5$ cm$^{-2}$. This falls within the error bounds of the column density derived through the relationship by Predehl & Schmitt [50] (1995):

$$N_{\text{HI}} = 1.79 \pm 0.03 A_V \times 10^{21} \text{cm}^{-2} \quad (14)$$

Using this relationship, a value of $N_{\text{HI}} = 22.13 \pm 0.5$ cm$^{-2}$ was obtained. Note that the neutral hydrogen mainly probes the region in front of the M17SW molecular cloud, where molecular $H_2$ concentrations are very low.
9 Spectral energy distribution

Information on the physical properties of the star ($L$, $T_{\text{eff}}$, and $R$) and the presence of circumstellar material can be derived from the spectral energy distribution (SED) of the source. First, the SED has to be corrected for interstellar extinction, using the empirically derived extinction law from Cardelli, Clayton & Mathis (1989) [42]:

$$A_\lambda/A_V = a(x) + b(x)/R_V$$

The $A_\lambda/A_V$ ratio (i.e. the absolute extinction) is a function of two wavelength dependent coefficients $a(x)$ and $b(x)$. With a choice of $R_V$ (defined in equation 15), the extinction in every filter can be calculated, if the extinction in the visual is estimated or known. The value of $R_V$ is a measure of the wavelength dependence of the extinction and is thought to be a measure of the dust grain size distribution. As $R_V$ gets larger, the extinction curve 'flattens' (see table 8).

<table>
<thead>
<tr>
<th>Filter</th>
<th>$\lambda$ ($\mu$m)</th>
<th>$m_{\text{obs}}$ (mag)</th>
<th>$A_\lambda/A_V$</th>
<th>$A_\lambda/A_V$</th>
<th>$A_\lambda/A_V$</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td></td>
<td></td>
<td>($R_V=3.1$)</td>
<td>($R_V=3.5$)</td>
<td>($R_V=3.9$)</td>
</tr>
<tr>
<td>U</td>
<td>0.36</td>
<td>17.50</td>
<td>1.56</td>
<td>1.50</td>
<td>1.44</td>
</tr>
<tr>
<td>B</td>
<td>0.44</td>
<td>17.02</td>
<td>1.31</td>
<td>1.30</td>
<td>1.27</td>
</tr>
<tr>
<td>V</td>
<td>0.55</td>
<td>15.55</td>
<td>1.00</td>
<td>1.00</td>
<td>1.00</td>
</tr>
<tr>
<td>R</td>
<td>0.70</td>
<td>13.60</td>
<td>0.75</td>
<td>0.76</td>
<td>0.78</td>
</tr>
<tr>
<td>I</td>
<td>0.90</td>
<td>12.21</td>
<td>0.48</td>
<td>0.50</td>
<td>0.52</td>
</tr>
<tr>
<td>J</td>
<td>1.25</td>
<td>10.59</td>
<td>0.28</td>
<td>0.30</td>
<td>0.30</td>
</tr>
<tr>
<td>H</td>
<td>1.65</td>
<td>9.29</td>
<td>0.18</td>
<td>0.20</td>
<td>0.21</td>
</tr>
<tr>
<td>K</td>
<td>2.2</td>
<td>8.05</td>
<td>0.11</td>
<td>0.12</td>
<td>0.12</td>
</tr>
</tbody>
</table>

Table 8: Calculated values of the $A_\lambda/A_V$ ratio as given in equation 15. UBVRI magnitudes are taken from Chini et al. (1980) [31], JHK from Hanson et al. (1997) [28].

The different values of $R_V$ were taken, because they represent the galactic total-to-selective extinction $R_V = 3.1$ and the value of $R_V$ is thought to increase with more developed dust environments, such as the M17 region. The $R_V = 3.9$ value is derived by Hoffmeister & Chini (2008) [32] (see section 4).

9.1 Photometric observations

With the corrected magnitudes $m_{\text{corr}}$, the spectral energy distribution (SED) of B275 can be reconstructed. In figure 19, $m_{\text{obs}}$ and $m_{\text{corr}}$ have been plotted on top of different Kurucz model atmospheres. The Kurucz 1993 catalogue of model atmospheres contains 7600 models, covering a large range of metallicities, gravities and effective temperatures.

A comparison between the theoretical parameters and the closest match from the Kurucz 1993 catalogue is given in table 9. The B2 V model was obtained by
interpolating between the B0 V and the B3 V model, as the Kurucz catalogue
did not offer a suggested model for a B2 V star. A model with $T_{\text{eff}} = 22000$ K
and $\log g = 4.0$ was taken for the B2 V spectral type.

<table>
<thead>
<tr>
<th>Spectral Type</th>
<th>$\log T_{\text{eff}}$ (K)</th>
<th>$\log T_{\text{eff,m}}$ (K)</th>
<th>$\log g$</th>
<th>$\log g_m$</th>
<th>$M$ (M$_\odot$)</th>
<th>$R$ (R$_\odot$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>O8 V</td>
<td>4.553</td>
<td>4.544</td>
<td>3.94</td>
<td>4.0</td>
<td>26</td>
<td>6.8</td>
</tr>
<tr>
<td>B0 V</td>
<td>4.477</td>
<td>4.477</td>
<td>3.9</td>
<td>4.0</td>
<td>15</td>
<td>5.0</td>
</tr>
<tr>
<td>B2 V</td>
<td>4.342</td>
<td>4.342</td>
<td>3.91</td>
<td>4.0</td>
<td>7</td>
<td>3.3</td>
</tr>
<tr>
<td>B5 V</td>
<td>4.188</td>
<td>4.176</td>
<td>4.04</td>
<td>4.0</td>
<td>4</td>
<td>2.7</td>
</tr>
</tbody>
</table>

Table 9: Parameters of the relevant Kurucz (1993) models, characterizing each
model atmosphere. The theoretical values of $T_{\text{eff}}$ and $\log g$ are given (taken from
Hanson et al. (1997) [28]), and the closest match from the Kurucz atlas (denoted
by the subscript m). The corresponding mass and radius for each spectral type
are taken from table 3 in Hanson et al. (1997) [28].

Figure 19 shows the discrepancy between the spectral type inferred from our
optical X-shooter spectrum and the spectral type from the extinction-corrected
SED. The optical classification yielded a spectral type of B6, but from the SED
fitting, a completely different conclusion has to be drawn. The 1300 pc models
reveal that the received flux from B275 is similar to that of an early B-star (this
is also shown in Hanson et al. (1997) [28]). With the distance of 2100 pc from
Hoffmeister & Chini (2008) [32], this discrepancy becomes even larger. B275
should then be classified as a late O type star.

The difference between the flux of a B5 and a B6 star is small, as the temper-

ature does not decrease as much per sub spectral type as with earlier type stars.
As a crude estimate, the SED should shift one unit down the logarithmic y-axis
(a factor of 10) to obtain the flux for the B5 V model (compared with the 1300
pc models). Then, to match the SED with the optical spectral classification,
M17 should be located at $\sim 412$ pc ($d = \sqrt{0.1 \cdot 1300^2} = 412$).

A similar argument can be given for the value of $A_V$. If we would have
overcorrected for extinction, the star would seem more luminous. A decrease
by a factor of 10 in flux corresponds with a decrease in magnitude of $\Delta m =
2.5$ mag, with $\Delta m$ the change in $A_V$, compared to the one used in figures 19.
According to this, the spectrum of B275 was overcorrected with 2.5 magnitudes
visual extinction, but the choice of $A_V$ (and $R_V$) is significant on the slope
of the extinction corrected SED. If B275 would be an OB star, an incorrect
combination of $A_V$ and $R_V$ would no longer reproduce the Rayleigh-Jeans tail
of the Kurucz model.

To conclude, the derived values for $A_V$ between 6.5 - 7.5 and $R_V$ between
3.1 - 3.5 seem to be accurate, as these values reproduce the Kurucz models well.
However, we note that there is a degeneracy between the $A_V$ and $R_V$, but based
on previous studies (e.g. Hanson et al. 1997, Hoffmeister et al. 2008), these
extinction parameters should apply to B275.
Figure 20: The observed magnitudes (square symbols) and the extinction corrected magnitudes (plus symbols) of B275, plotted on top of several Kurucz (1993) model atmospheres for ZAMS OB stars, with different values of $R_V$ (3.1, 3.5, 3.9). The left panels place the model at a distance of 1300 pc, while the right panels assume a 2100 pc distance. The corrected magnitudes are plotted with different amounts of extinction ($A_V = 6.5, 7, 7.5$ mag). Note the infrared-excess, appearing at $\sim 1 \mu m$, where the radiation from the circumstellar material (from 1500 K) starts to contribute. The surrounding material of B275 will be discussed in section 10).
As can be seen in figure 19, a spectral type of O9-B2 would fit the SED of B275 at a distance of 1300 pc. O8-B1 would be the spectral type of choice when M17 is located at 2100 pc. The luminosity of a star is related to the bolometric magnitude according to:

$$\log\left(\frac{L}{L_\odot}\right) = -0.4(M_V + BC - M_{bol,\odot})$$

Where the bolometric magnitude of the sun is taken to be $M_{bol,\odot} = 4.75$ (Allen, 1976 [43]). Bolometric magnitudes were taken from the Geneva models as a function of effective temperature published in Schaller et al (1992) [44]. As a result, the photometric flux would exceed the flux through optical classification by a factor of 400 in the most extreme scenario.

### 9.1.1 Comparison with mYSOs inside NGC6618

The massive YSOs from Hanson et al. (1997) [28] in the direct vicinity of B275 were compared with B275, in order to search for a systematic discrepancy between luminosity and optical classification in objects located in M17 (see table 11). Even though there are stars with similar characteristics as B275 (CO first overtone, Pa$\delta$, similar $A_V$), no other star listed in Hanson et al. (1997) [28] contains such an obvious discrepancy as B275.

### Table 10: Bolometric magnitudes and calculated luminosities for the used spectral types.

<table>
<thead>
<tr>
<th>Spectral type</th>
<th>$M_{bol}$</th>
<th>$\log\left(\frac{L}{L_\odot}\right)$</th>
</tr>
</thead>
<tbody>
<tr>
<td>O8 V</td>
<td>-7.58</td>
<td>4.932</td>
</tr>
<tr>
<td>O9V</td>
<td>-7.10</td>
<td>4.740</td>
</tr>
<tr>
<td>B1V</td>
<td>-4.63</td>
<td>3.752</td>
</tr>
<tr>
<td>B2 V</td>
<td>-3.59</td>
<td>3.336</td>
</tr>
<tr>
<td>B6 V</td>
<td>-0.90</td>
<td>2.260</td>
</tr>
</tbody>
</table>

Table 11: Comparison of the massive YSOs in M17. The sample was chosen from the stars with plotted SEDs in Hanson et al. (1997) [28], so the optical classification could be compared with an extinction corrected SED. Some spectral features are listed in the table to extend the comparison. Several sources show double peaked emission (DPE), other single peaked (E).

<table>
<thead>
<tr>
<th>Name</th>
<th>Sp. T.</th>
<th>SED</th>
<th>Pa$\delta$</th>
<th>CO(2-0)</th>
<th>$A_V$</th>
</tr>
</thead>
<tbody>
<tr>
<td>B275</td>
<td>B6 V</td>
<td>Late O - Early B</td>
<td>DPE</td>
<td>Strong</td>
<td>7.5</td>
</tr>
<tr>
<td>B243</td>
<td>Early B</td>
<td>Early B</td>
<td>DPE</td>
<td>Weak</td>
<td>8.6</td>
</tr>
<tr>
<td>B331</td>
<td>-</td>
<td>Late O - Mid B</td>
<td>DPE</td>
<td>Yes</td>
<td>$\sim 12$</td>
</tr>
<tr>
<td>B268</td>
<td>B2 V</td>
<td>Early B</td>
<td>E</td>
<td>Yes</td>
<td>7.8</td>
</tr>
<tr>
<td>B260</td>
<td>O7-O8</td>
<td>O7</td>
<td>-</td>
<td>-</td>
<td>8</td>
</tr>
<tr>
<td>B289</td>
<td>O9.5V</td>
<td>Late O - Early B</td>
<td>-</td>
<td>-</td>
<td>7.7</td>
</tr>
<tr>
<td>B337</td>
<td>-</td>
<td>Early B</td>
<td>-</td>
<td>Strong</td>
<td>11.5</td>
</tr>
</tbody>
</table>

Table 11: Comparison of the massive YSOs in M17. The sample was chosen from the stars with plotted SEDs in Hanson et al. (1997) [28], so the optical classification could be compared with an extinction corrected SED. Some spectral features are listed in the table to extend the comparison. Several sources show double peaked emission (DPE), other single peaked (E).
9.2 Flux calibrated spectrum

The fact that the luminosity of B275 far exceeds the ZAMS value, corresponding to the temperature of the photosphere, indicates a larger radius. This would confirm giant features of the spectral lines, seen in section 7.2. By increasing the emitting surface of the star, the radius can be calculated that would correspond with the observed flux (Chini (1980) [31]). This was done by de-reddening the flux calibrated spectrum, using the extinction law from CCM. The continuum level at the interval 400-800 nm was used to represent a clean photospheric spectrum slope, without the contribution of reprocessed stellar radiation from the circumstellar material. Using a chi-square test, the A_V and R_V for the 400 - 800 nm region were determined, reproducing the slope of the theoretical Kurucz model. R_V = 3.3 and A_V = 6.1 were derived, which correspond well with the parameters derived in section 9.1. The flux scales with the factor (R/d)^2, where R is the radius and d is the distance towards the object. This dimensionless value was used to derive the corresponding radius at a distance of 1300 pc and 2100 pc, respectively. An example is drawn in figure 21. For the Kurucz model, a B7 III model atmosphere (T_eff = 13,000 K; log g = 3.5) was used, because the height of the Balmer jump was best reproduced with a model of this temperature.

![Example of fitting the observed de-reddened spectrum (orange) with a theoretical Kurucz model (blue). On the left, M17 is located at 1300 pc, while on the right the radius is calculated if M17 would be located at 2100 pc. Best fit extinction parameters are R_V = 3.3 and A_V = 6.1. The red squares illustrate the photometric observations, taken from Chini et al. (1980) [31]). Merging problems arise between the different X-shooter orders in the blue part of the spectrum.](image)

Figure 21: Example of fitting the observed de-reddened spectrum (orange) with a theoretical Kurucz model (blue). On the left, M17 is located at 1300 pc, while on the right the radius is calculated if M17 would be located at 2100 pc. Best fit extinction parameters are R_V = 3.3 and A_V = 6.1. The red squares illustrate the photometric observations, taken from Chini et al. (1980) [31]). Merging problems arise between the different X-shooter orders in the blue part of the spectrum.

The ZAMS radius of a B7 V star is ∼ 2.5 R_⊙, while a B7 III star would have a radius of ∼ 7 R_⊙ (Schmidt-Kaler [46] (1982)).
We derived $5.3 \, R_\odot > R > 8.6 \, R_\odot$ for B275. At a distance of 1.98 kpc (Xu et al. (2011) [33]), B275 has a radius of $8.1 \, R_\odot$. This value is consistent with our observations of the Balmer line widths, explained in section 7.2.

### 9.3 HR-diagram

With the $T_{\text{eff}}$ and $L$ (and thus: $R$), the location of B275 in the HR-diagram is well constrained.

Figure 22: Left: The two upper solid curves represent the stellar birthlines for intermediate mass stars for a constant mass accretion rate of $\dot{M} = 10^{-5} \, M_\odot/\text{yr}$ and $\dot{M} = 10^{-4} \, M_\odot/\text{yr}$ calculated by Palla & Stahler (1992) [7] (see section 3). The dashed lines originating from the birthline are PMS tracks according to their stellar mass (1-6 $M_\odot$). The dotted lines represent lines of constant stellar radius of 1 $R_\odot$ (lower) and 10 $R_\odot$ (upper). Right: Plotted are accreting protostar tracks, calculated by Hosokawa (2010) [16]. The black squares represent the OB stars in M17 and B275 is bound by the red box. The dashed line inside the box shows the position of B275 if M17 were located at 1.98 kpc (Xu et al. 2011 [33]).

Figure 22 shows accreting protostar and PMS tracks. The dotted lines are lines of constant stellar radius and were calculated by assuming that the stars radiate as black bodies:

$$\left(\frac{L}{L_\odot}\right) = \left(\frac{R}{R_\odot}\right)^2 \left(\frac{T}{T_\odot}\right)^4$$

The location of B275 is bound by the uncertainty in spectral classification (B6 - B7) and the uncertainty in distance towards M17, which led to a range in radii calculated in section 9.2.
10 Circumstellar environment

The position of B275 in figure 7 and the infrared excess, seen in figures 19 and 21, have confirmed the presence of circumstellar material around B275. The K-band excess for B275 is \( \sim 3 \) mag, because the J-K or H-K color of a B6 star is \( \sim 0 \) mag (Hanson et al. (1997) [28]). In the introduction of this thesis, it was shown that radiation of a massive object (\( \sim 8 \, M_\odot \)) may halt the infalling material, preventing further growth of the central object. However, there is growing evidence that massive stars do form by the accretion of gas through a disk like low-mass stars (Keto & Wood (2006) [4], Zinnecker & Yorke (2007) [18]). The formation of a massive accretion disk could be the key in coping with the devastating radiation pressure of the mYSO. In this section, the observed properties of the material surrounding B275 will be discussed.

10.1 Evidence for a Keplerian disk

The optical spectrum clearly shows the near-infrared Ca ii \( (8498, 8542, 8662 \, \text{Å}) \) triplet, the O i 8446 Å and possibly the H\( \alpha \) and H\( \beta \) lines with double peaked emission, as seen in figures 25 and 26. The higher transitions of Balmer, Paschen and Brackett series also show emission features.

![Figure 23: The origin of the radiation from a Keplerian emission line. Figure adapted from Isella et al. (2007) [52].](image)

Figure 23 shows how a double peaked Keplerian profile is produced. Point A, at the central wavelength, originates from the region directly in front or behind the object and does not show Dopplershift. Regions B and C comprise the largest part of the profile and contain the largest amount of flux. These regions do show Dopplershift, as the disk rotates. Since a normal Keplerian disk has a velocity structure with \( v \propto \sqrt{1/r} \), with \( r \) the radius of the disk, the highest radial velocities will occur near the central object. These velocities correspond with the line wings of the Keplerian profile (point D).
Figure 24: The hydrogen Paschen 13, 15 and 16 lines are filled in with the Ca II triplet. The Ca II and O I 8447 lines are double peaked, which could reveal their circumstellar nature. Central emission components are visible in the Pa 12 an 14 lines.

CO first and second overtone emission at 2.3 µm and 1.5 µm are also detected, which is indicative of a dense and shielded region. This is because CO is easily dissociated by UV radiation, but also requires a high temperature for excitation. The shape of these emission features reveal that they likely originate from a circumstellar disk. The CO overtone emission will be discussed in section 10.2.6.

10.2 Structure of the disk and constraints on emitting regions

Even though line profiles and excitation conditions of the observed transitions can give detailed information on the structure and composition of the circumstellar material, the exact origin of emission lines in mYSOs is still not known. Various models and mechanisms are proposed to validate the observations, which could give us more insight in the formation process of the most massive stars.

The peak separation in Keplerian profiles is approximately half of the projected velocity at the outermost radius of the emitting region (Carr 1995 [53]). A large peak separation could indicate that the emission line traces the inner disk, where the rotational velocity is high. By comparing the peak separation of emission profiles with information on excitation conditions, we can constrain the region where the emission line is produced.

The Ca II triplets and the hydrogen lines Hα and Hβ have comparable peak separation (Δv ~ 105 km s⁻¹), while O I shows a less separated structure (Δv ~ 70 km s⁻¹). The excitation energy of Ca II is small (~ 3.1 eV) compared to O I 8447 (~ 10.9 eV).
The strength of an emission line depends on four main factors:

1) The optical depth of the line emitting region. As long as the line is optically thin, all material can contribute to the line emission.

2) The amount of material.

3) The size of the line emitting region.

4) The value of the Einstein coefficient $A_{ij}$ for spontaneous emission.

As a consequence, the region inside the circumstellar disk where the emission originates from, depends on a combination e.g. density, abundance, temperature, line strength and the size of the line forming region.

10.2.1 The ionized surface layer: hydrogen lines

The hydrogen recombination lines originate from the regions where the hydrogen gas is ionized, i.e. locations accessible to photons with energies $E \geq 13.6$ eV. In the upper layers of a circumstellar disk, where the density is relatively low, the ionizing radiation from the central star can still penetrate the material.

Shown in figure 26 are the Balmer, Paschen and Brackett series of hydrogen. The increase in strength of the central emission components of hydrogen through the hydrogen series can be explained as follows: as soon as hydrogen gets ionized by the central (proto)star, an electron and a proton recombine in any energy level. The electron subsequently cascades down to the ground level, emitting radiation at any transition. The lower transitions will more frequently take part in such a cascade, resulting in a greater line strength. In addition, the Einstein $A_{ij}$ coefficient is larger for the lower hydrogen transitions. The hydrogen lines are
often used to estimate the density of the emitting gas, because weaker lines of hydrogen require a higher electron density compared to the stronger line series of hydrogen (Balmer, Paschen and Brackett series).

The optical depth of the emitting gas affects the relative line strengths of hydrogen. In low density regions, the strength of the hydrogen lines can be calculated based on the case B approximation, where the ratios of the hydrogen lines are well determined (Storey & Hummer (1995) [54]). In high density regions, where the gas is optically thick, the strength should be determined by the emitting surface of the gas. Exact ratios in the spectrum of B275 are hidden from view by the absorption lines from the stellar photosphere and need further investigation.

The double peaked profiles of Hα and Hβ are surprising, because these are the strongest transitions of hydrogen and, therefore, are expected to originate from large extends in the disk. This would result in a single peaked emission feature, whereas the higher transition of hydrogen would be formed relative close to the star, where the profiles are rotationally broadened. These transitions would show double peaked emission more frequently. The double peaked emission in the strong hydrogen lines could be due to an artifact of the reduction process, or wrongly subtracted nebular contamination.

Figure 27: Hβ profile. The X-shooter spectrum (black) is fitted with a Voigt profile (red). The bisector (purple) of the photospheric profile is calculated and after subtracting the fitted profile the resulting spectrum (blue) is plotted. In addition, the normalized Hα emission profile (green) is shown for comparison. The dotted lines (the width of the Hα emission) mark the region excluded from the fit.

We also noted a mismatch in the Hβ profile shown in figure 27, as the emission component is not located centrally in the absorption profile. This trend was
also seen in several Paschen profiles, most notably Pa-6 and Pa-7 (see figure 26). The feature seen in the red wing of the Hβ is the 4882 Å diffuse interstellar band and the measured shift of the photospheric line $v = 46.6 \text{ km s}^{-1}$ is not compatible with measurements of the system velocity (centre of double peaked structures and other photospheric lines) $v_{\text{system}} \sim 20 \text{ km s}^{-1}$.

10.2.2 H ii region: Nebular lines

The emission profiles from B275, which is located within a H ii region, could be contaminated with recombinations from the gas in the ionized region. The high energetic photons from the star ionize the gas in the H ii region, followed by a recombination. The recombination will be visible as a sharp emission peak, because the gas is not rotationally broadened inside the H ii region, in contrast to the upper layer of the circumstellar disk. The FHWM of the nebular emission is close to the sound speed in ionized gas ($\sim 10 \text{ km s}^{-1}$), for the gas originating from the disk this could be significantly higher (the central emission components of the hydrogen recombination lines show profiles 100-150 km s$^{-1}$, see figure 26). Nebular lines can also be distinguished from spectral lines by looking at a two dimensional spectrum: the nebular lines will be emitted on the whole spatial range of the slit, the spectral lines only on the position of the object. The X-shooter pipeline normally subtracts the nebular emission during the reduction of the science frames. In our spectrum, Hα and Hβ could be suffering from nebular contamination, where the nebular emission from these two lines has not subtracted correctly, resulting in a sharp double peaked profile. However, the nebular spectrum of B275 needs more investigation in order to determine its influence on the derived properties of the circumstellar features.

10.2.3 [OI]

Forbidden transitions de-excite very slowly and this is why they must come from low density regions, where the collisional rates are suppressed sufficiently for particles to radiate in forbidden line transitions.

Störzer & Hollenbach (2000) [55] proposed two different mechanisms, which contribute to the line strength of forbidden oxygen:

1. In high temperature regions ($T > 3000 \text{ K}$), the oxygen atoms can be thermally excited by hydrogen atoms and free electrons.

2. The excited oxygen atoms can be produced after photodissociation of OH by UV radiation.

The detected [O i] 6300 Å and 6363 Å forbidden transitions (figure 29) are not resolved and are detected along the whole slit. The lines are also seen in the sky spectrum of the Earth’s atmosphere and therefore, no conclusion can be drawn about their origin.

Figure 28: The forbidden oxygen line at 6300 Å.
10.2.4 The Ca \textsc{ii} triplet

The Ca \textsc{ii} triplet lines arise from the same upper level (3P\textsuperscript{6}4P) as the calcium H and K lines at 3969 and 3934 Å. The lower levels of the triplet transitions are metastable and decay in forbidden transitions (7291 and 7324 Å) to the Ca \textsc{ii} ground state 3P\textsuperscript{6}4S at a rate of \( \sim 1.3 \) s\(^{-1}\). The absence of the [Ca \textsc{ii}] lines in our observed spectrum reveals that the Ca \textsc{ii} emission comes from the dense circumstellar environment, where the forbidden transitions are collisionally de-excited before they can radiate to the ground state through the forbidden line transition. In our case, the triplet features are optically thick, as the peak ratios of the several components are more close to unity than the 1:9:5 optically thin values (derived from their \( g_u A_u h \nu \) values [47]). As calcium is not very abundant, the region where the Ca \textsc{ii} triplet is to be emitted might lie where most of the calcium is singly ionized. Ca \textsc{iii} appears rapidly where hydrogen is ionized, because the ionization potential of Ca \textsc{ii} is 11.8 eV. The Ca \textsc{ii} triplet region should therefore be shielded from photons with energies \( E \geq 13.6 \) eV. The ionizing flux decreases inside the disk and as soon as the photons of \( E \geq 11.8 \) eV are sufficiently absorbed, the Ca \textsc{ii}/Ca \textsc{iii} ratio will increase. Neutral carbon is an effective element to protect the calcium atoms from the destructive 11.8 - 13.6 eV radiation coming from the star. It has an ionization potential of 11.26 eV and a much larger abundance \( [n(C)/n(Ca)] \) with a larger photoionization cross section. Persson (1988) [56] showed that the Ca \textsc{ii}/Ca \textsc{i} fraction is high for a large range in electron densities at 3000-4000 K.

10.2.5 O \textsc{i} 8446 Å emission

The O \textsc{i} 8446 Å excitation requires the oxygen to be predominantly neutral. Since there exists a very efficient charge exchange between O \textsc{i} and H \textsc{i} (O \textsc{i} + H \textsc{i} \rightarrow O \textsc{ii} + H \textsc{i}) this must be located where hydrogen is also neutral. The O \textsc{i} 8446 line can be excited in three different ways. Firstly, in pure recombination, the 7774 Å line should be present [56]. In our spectrum the 7774 O \textsc{i} feature is in absorption, so we can rule out recombination as the main component of the 8446 Å line formation. The second and third option is that the excitation is caused by fluorescence or Ly\( \beta \) emission. In case of continuum fluorescence, the subsequent cascade emits the 13164 Å line, which is not detected. In addition, the 11287 Å line is emitted, which could not be detected because of strong telluric absorption at this wavelength. The Ly\( \beta \) flux is easily absorbed by hydrogen atoms and therefore it is not likely that these high energy photons will be able to reach
the neutral oxygen. As a result, the origin of the [O i] 8446 Å emission remains inconclusive.

10.2.6 The dense inner disk: CO bandheads

At the innermost region of a circumstellar disk, the local density can be high enough for molecules to survive. The 2.3 μm CO (2-0) emission bandhead profile is an example of such a feature. Only at high temperatures the CO molecule is sufficiently excited to emit a ro-vibrational spectrum in the near-infrared. Therefore, it has been proposed to originate from the hot and dense inner part of the circumstellar disk of a YSO ($N_H \approx 10^{10}$, $1500 > T > 4500$ K) and is thus a powerful diagnostic of disc material close to the stellar surface (e.g. Wheelwright et al. (2010) [57]). The emitting region is located within a few AU of the central star, with a maximum distance depending on the inclination of the system. When the density of the disk is sufficiently high ($10^{20} - 10^{21}$ cm$^{-2}$), the molecules are able to withstand photodissociation from the central star through self shielding. In addition, CO molecules can form within the gas phase to compensate for destruction or dissociation. Bik & Thi (2004) [64] showed that a lot of CO bandheads are fitted well by an optically thin Keplerian disk model. The presence of a 'blue shoulder' in the CO emission profiles, which is detected in the spectrum of B275, is characteristic of a large inclination. The detected CO emission profiles of B275 are shown in figure ???. We also note the presence of the CO (3-0) second overtone emission at 1.5 μm in the spectrum of B275. This transition has an even higher excitation temperature and, as a result, is expected to originate even closer to the star compared with the CO first overtone profile. However, no significant rotational broadening is detected in the CO second overtone profiles.

![Figure 30: Left: the CO(2-0) to CO(5-3) (black) and CO(3-0) to CO(6-3) (blue) bandhead emission features. Right: Same as the upper panel, but smoothed with 7 pixels to minimize noise.](attachment:figure30.png)
10.2.7 Emitting regions in the disk

After the discussion on the different elements present in the NIR spectrum of B275, we can summarize our findings using a schematic of the circumstellar environment, shown in figure 31.

Figure 31: A schematic view of a circumstellar disk around a mYSO (figure adopted from Bik 2008). Plotted inside are the emitting regions of the discussed elements.

Figure 31 shows a possible cross-section of a disk around a mYSO. The possible emitting regions of the elements are drawn inside, based on the findings in section 10.2. The region closest to the star is gaseous, as dust can only be present outside the dust sublimation zone at $\sim 1500$ K (however, this depends on the amount of self shielding). We see the star-disk system B275 under a sufficient large angle which allows the direct detection of the photosphere. The inclination could be constrained by modeling the CO bandhead profiles (Bik & Thi 2004 [64]). The hydrogen lines are created in the ionized upper layer of the disk and possibly a disk wind. The enormous strength of the lower hydrogen lines are partly explained by a larger emitting surface area. A disk wind results in a radiation driven outflow with velocities up to 200 km s$^{-1}$ (Drew et al. 1998 [60]), which is compatible with the widths measured in the emission profiles of B275. Forbidden line transitions are also often seen in a disk wind, but as the line is not resolved, the origin of the [O i] line can not be constrained. The Ca II triplet originates from a warm ($\sim 3000 - 4000$ K), shielded region for enough singly ionized calcium particles to survive (Persson 1998 [56]). The region where a significant fraction of calcium is singly ionized is located in the inner parts of the circumstellar disk. These conditions are similar to that of the CO overtone features. The O i line forming mechanisms are discussed, but the dominant mechanism of O i emission in B275 is unclear. As a general remark, we note that the O i emitting region could be extending towards a larger radius as Ca II, because O i shows a smaller velocity separation in its emission profile.
11 Discussion

After the analysis of the X-shooter spectrum of B275, we have come to the following conclusions:

I) Optical classification yielded a spectral type B5 III - B6 III. With the flux calibrated spectrum and models generated with FASTWIND, B7 III was concluded.

II) The discrepancy between photospheric spectral type and luminosity can be explained by allowing for a larger radius. These results are consistent with the trend seen in the Balmer line widths. With the recent measured distance of $1.98^{+0.14}_{-0.12} \text{kpc}$ (Xu et al. (2011)) and a B7 III model atmosphere (Kurucz (1993)), a radius of $8.1 \, R_\odot$ was determined for B275.

III) According to its position in the HR diagram, B275 is a pre-main sequence star still heading towards the main sequence.

IV) Although no direct evidence for ongoing accretion has been seen (e.g. UV and optical excess flux), B275 is still surrounded by a dense Keplerian disk close to the star, as indicated by the double-peaked emission profiles of the calcium triplet and the presence of the CO first and second overtone emission. Also, no outflow has been detected by jets or shocks, which could be an inclination or extinction effect, or due to a temporary halt in the accretion process.

V) B275 is compatible with accretion tracks of Palla & Stahler (1992) and Hosokawa (2010). If we assume the accretion to be halted, B275 does not lie on a PMS track of $10^{-5} \, M_\odot \, \text{yr}^{-1}$. According to their models, increasing the preceding accretion rate to $\sim 10^{-4} \, M_\odot \, \text{yr}^{-1}$ would shift the complete stellar birthline to higher mass. In this case, B275 would have a defined PMS track and would be contracting on a Kelvin-Helmholtz timescale, but inhomogeneities in the accretion material could lead to strong variations in accretion diagnostics.

VI) The DIBs in M17 do not correlate with the line-of-sight extinction, which was also shown by Hanson et al. (1997). Possible explanations include that radiation destroys the DIB carrier or the DIB features get saturated after a certain column density of carriers. Hoffmeister et al. (2008) showed that the total extinction towards B275 seems to be composed of a galactic foreground value ($R_V = 3.1$), plus an intrinsic ($R_V = 3.9$) value. We have derived $R_V = 3.3$ towards B275.

After the spectral classification, discussed in section 7, and comparing the outcome with the observed photometric measurements in section 9, it became clear that the luminosity of B275 exceeds the luminosity of a main sequence B6 star by a factor of 400. Indications for the extra luminosity came from the width of the Balmer lines, shown in section 7.2, which showed signatures of decreased
pressure in the photosphere and caused the line widths to be smaller compared to a main sequence star. FASTWIND models (Puls et al. 2005 [45]) yielded a best fit with a B7 III model ($T_{\text{eff}} = 13,000$ K; log $g = 3.5$) and showed that the radius of B275 exceeds its main sequence value and, subsequently, the radius of B275 was determined by fitting the flux calibrated spectrum on a B7 III Kurucz model (section 9.2). A radius of $R = 8.1 \, M_{\odot}$ was determined for B275, whereas an B6 V star would have a radius of $R = 2.5 \, M_{\odot}$ (Schmidt-Kaler (1982) [46]). As a result, B275 was placed in the HR-diagram and the pre-main sequence nature of the object was concluded. An estimation of the current mass of B275 can be made using:

$$\left( \frac{M}{M_{\odot}} \right) = \left( \frac{g}{g_{\odot}} \right) \left( \frac{R}{R_{\odot}} \right)^2$$

Adopting a surface gravity for B275 of log $g = 3.5$ (used in the FASTWIND models) and log $g = 4.438$ for the Sun (Allen 1976 [43]), the current mass of B275 is estimated at $M \approx 7.5 \, M_{\odot}$.

The position of B275 in the HR-diagram is compatible with the models of Palla & Stahler (1992) [7]. According to their computations, the evolutionary status of B275 can be explained as follows:

1) B275 has finished accretion and is gravitationally contracting on a Kelvin-Helmholtz timescale. In order to estimate the duration of this timescale, we use a mass of $M = 7.5 \, M_{\odot}$ (equation 18) and with equation 8, along with typical numbers for the luminosity for a O8 - B2 star, we can estimate the remaining KH-timescale to be of the order $10^3 - 10^4 \, M_{\odot} \, \text{yr}^{-1}$. As can be seen in figure 22, B275 is located well above the birthline calculated for $\dot{M} = 10^{-5} \, M_{\odot} \, \text{yr}^{-1}$. Palla & Stahler (1992) [7] showed that higher accretion rates leads to a shift of the birthline in the HR-diagram (see figure 4). B275 does have a PMS phase if we assume that the star accreted material at a rate of $\dot{M} = 10^{-4} \, M_{\odot} \, \text{yr}^{-1}$ in the past.

2) B275 lies on proto stellar accretion tracks with mass accretion rates $10^{-5} \, M_{\odot} \, \text{yr}^{-1} < \dot{M} < 10^{-4} \, M_{\odot} \, \text{yr}^{-1}$ calculated by Palla & Stahler (1992) [7] and Hosokawa et al. (2010) [16]. The models of Hosokawa (2010) [16] have shown that a massive star indeed significantly ‘swells’ during disk accretion and subsequently contracts on a KH-timescale. The absence of direct accretion signatures as optical veiling can be explained through inhomogenities in the accretion flow.

The main difference in these scenarios is whether B275 has crossed the birthline, i.e. the location in the HR diagram where accretion stops and the object becomes optically visible (described in section 3). However, one can imagine that, in practice, an abrupt cease of accretion is not likely in the case of B275, where we have detected a dense inner molecular disk. While the infrared excess, double-peaked lines profiles (Hα, Ca II and O i) and CO first and second overtone bandheads indicate the presence of circumstellar material close to the star,
no evidence for ongoing accretion has been detected through UV and optical excess flux, inverse P-Cygni profiles or jet/outflow signatures (Ghandour et al. 1994 [61]). Other accretion diagnostics, such as the Ca II triplet and Hα profile, could possibly be used to constrain the current mass accretion rate, as has been shown in lower mass stars (Muzerolle et al. 1998 [62]). However, the accretion process in massive YSOs is until now poorly understood, in particular because of the strong variation on small timescales (R. Chini, private communication). Future observations could possibly be used to determine whether B275 is still accreting or not and, furthermore, large field-of-view images of the M17 region could reveal outflow signatures from B275, hidden from the spectroscopic slit of X-shooter. Nonetheless, if we assume B275 is gravitationally contracting with constant L towards the ZAMS, the final mass can be estimated at \( M \sim 7.5 \, M_\odot \), as given by equation 18. This would represent a lower limit to the final mass.

One may argue that the spectrum of B275 bears similarities to that of a classical Be star or Be shell star. However, only B[e]-supergiant stars are suffering from such heavy mass loss that is sufficiently dense to emit CO first and second overtone emission (e.g. McGregor et al. 1988 [68]). In addition, Be stars would rotate at or near the critical break-up speed, which would be reflected by the same rotational speed of the star and the gas at the stellar surface, i.e. \( v \sin i = v_{\text{gas}} \sin i \). However, as the maximum speed of the Ca II profile is comparable with the rotational velocity of the star \( (v \sin i = 115 \pm 11 \, \text{km s}^{-1}) \), means that there is gas close to the star which is rotating faster than the stellar surface, i.e. \( v \sin i < v_{\text{gas}} \sin i \). This would not be possible in case of mass loss at the equator and, consequently, we can exclude B275 to be a classical B emission star.

The SED is typical of Herbig Be stars, showing an infrared excess, starting at 1 micron, but Herbig Be stars often appear rather isolated, relatively close by and possess very long lived disks (Waters & Waelkens 1998 [65]). B275 is different, because of its high location in the HR-diagram and the fact that it is located in a massive star forming cluster. The disk around B275 could be a remnant of a previously massive and extended disk, where the outer parts have already been photo-evaporated by the presence of a strong stellar wind (Shepherd 2001 [69]). The actual geometry, composition and evolutionary status of the disk can possibly by explained by longer wavelength studies and the detection of accretion signatures would reveal the disk to be active or passive. Concluding, valuable information can be extracting from further research on the disk.

The mismatch in the absorption and emission of the Hβ profile could be indicative of a complex infall geometry, where part of the redshifted emission is absorbed by infalling material in front of B275, or that B275 is part of a binary system. However, the mismatch could also be an artifact of an incorrect nebular subtraction during the reduction of the data. A satisfying explanation of the Hβ profile has, until this point, not yet been found.

As a final remark, we note that B275 is located far from the main sequence compared with the massive stars in M17. These OB stars have most likely left the main sequence, due to their short lifetimes. As can be seen in figure 8,
the distribution of massive stars systematically differ from the mYSOs, which are located inside the M17SW molecular cloud. Either the massive stars have triggered the onset of star formation in the M17SW region, or B275 is part of the massive cluster. This can provide information on the age of the NGC6618 cluster, where B275 could be the next star reaching the main sequence.
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