Chapter 1

Introduction

At night, away from the nocturnal glow of urbanised areas, a diffuse band of light emerging from a myriad of stars can be discerned, reaching from horizon to horizon. These stars constitute the stellar disk of our Galaxy, the Milky Way. The space between these stars is vast, but it is far from empty. Everything in between the stars is called the Interstellar Medium (ISM). A tenuous gas, filling most of interstellar space, is structured in distinct phases. Three main phases exist in approximate pressure ($P$) equilibrium with each other, i.e., the product of the density ($n$) and temperature ($T$) in each phase is constant ($P \approx nT \approx 10^4 \text{ cm}^{-3} \text{ K}$).

These are the hot coronal and intercloud gas ($n < 10^{-2} \text{ cm}^{-3}, T \sim 10^6 \text{ K}$); the warm, neutral and ionised gas ($n \sim 0.1-1 \text{ cm}^{-3}, T \sim 10^3-10^4 \text{ K}$); and the cool atomic gas ($n \sim 100 \text{ cm}^{-3}, T \sim 100 \text{ K}$). An additional hot, ionised gas phase ($n \sim 1-10^5 \text{ cm}^{-3}, T \sim 10^4 \text{ K}$) is present in the form of discrete nebulae called HII regions. In both mass and volume, they occupy a negligible fraction ($\leq 1\%$) of the ISM, but they are very prominently visible throughout the Galaxy. In contrast to the previous phases, HII regions are not in equilibrium but continuously fed by stellar activity.

The final phase comprises the cold molecular gas ($n \geq 200 \text{ cm}^{-3}, T \sim 10 \text{ K}$) that is confined to giant molecular clouds. The clouds are held together by self-gravity rather than being in equilibrium with the other phases. Molecular clouds represent the locations in the Galaxy where the star-formation process is initiated.

Even to the naked eye, some nearby molecular clouds are visible in silhouette against the bright background of the Milky Way. These regions were first described as being ‘holes in the sky’ by Sir William Herschel (1785). This perspective did not change until Edward E. Barnard recognised in a photographic survey (1919) that the dark regions were actually discrete nebular objects blocking out the light from more distant stars. Halfway through the 20th century the association of dark nebulae with star formation became apparent. Bart J. Bok (1948) claimed that the dark nebulae, now called ‘Bok globules’, are objects where new stars are formed.

---

Remarkably, Emanuel Swedenborg proposed the nebular hypothesis much earlier (1734), based on philosophical arguments. The philosopher Immanuel Kant developed this idea further in 1755 and reasoned that the Solar System originated from a slowly rotating, collapsing nebula.
Present-day technology in infrared and (sub-) millimetre observations allow the direct detection of embedded, low-mass star formation in nearby dark clouds. The term ‘low-mass’ refers to stars with a mass similar to, or less than, about one solar mass \((1 \, M_\odot \approx 2 \times 10^{33} \text{ g})\). ‘Nearby’ in this respect means within a few hundred parsec \((1 \, \text{pc} \approx 3.26 \text{ lightyear} \approx 3 \times 10^{16} \text{ m})\). Stars more massive than \(8 \, M_\odot\) are referred to as high-mass, or massive stars. The recent discovery of distant \((\gtrsim 1 \, \text{kpc})\) and hence more massive dark cloud complexes, the so-called Infrared Dark Clouds, provides the opportunity to study the origin of massive stars and stellar clusters.

This thesis presents a study of the physical characteristics and star-forming properties of dark and massive molecular clouds throughout the Galaxy, with a strong emphasis on the existence of these clouds in the Outer Galaxy. The first chapter introduces various aspects related to the process of star formation. After that, several strategies for observing the early stages of star formation are alluded to. In the final section, an outline of the remaining chapters is given.

1.1 From clouds to stars: A general overview

The cold molecular phase of the ISM consists of a hierarchical structure of molecular clouds, from the giant cloud complexes to the small, gravitationally unstable cores. These different appearances are primarily described by their diversity in size, mass and density. A spontaneous or triggered collapse of the densest parts of molecular clouds initiates the process of star formation.

The first part of this section gives a qualitative description of the observed properties of different manifestations of molecular ‘clouds’ in the Galaxy. The use of a consistent nomenclature is somewhat ambiguous, not in the least because of the wide variety in properties. Moreover, in the astronomical literature there is no clear consensus so far on a uniform terminology. Nonetheless, throughout this thesis I will use the descriptions given in Section 1.1.1 whenever possible.

The second part of this section gives an introduction to the relevant time-scales during star formation and summarises our current understanding of the different stages of both isolated low-mass star formation as well as the formation of stellar clusters. The typical shape of the stellar mass-spectrum, referred to as the initial mass function (IMF), is briefly mentioned at the end of the section. Extensive sources of information on both observational and theoretical issues in star formation are, e.g., Stahler & Palla (2005) and McKee & Ostriker (2007). Most of the information in this chapter is adapted from these references and additionally from Bergin & Tafalla (2007) and Tielens (2005), unless specifically stated otherwise.

1.1.1 The hierarchical structure of molecular clouds

Clouds

*Molecular Clouds* (MCs) constitute most of the cold molecular gas in the Galaxy. They span a wide range of sizes, from small molecular clouds with sizes \(\sim 20 \, \text{pc}\) and masses \(\sim 10^4 \, M_\odot\) (e.g., Taurus MC) to the giant molecular clouds (GMCs) and complexes of sizes \(\sim 100 \, \text{pc}\) and masses \(\sim 10^5–10^6 \, M_\odot\) (e.g., the Orion MC). A typical gas
Table 1.1: Overview of the global properties of molecular clouds.

<table>
<thead>
<tr>
<th>Cloud Type</th>
<th>MCa</th>
<th>IRDC</th>
<th>Clumps</th>
<th>Cores</th>
<th>Globules</th>
</tr>
</thead>
<tbody>
<tr>
<td>size (pc)</td>
<td>20–100</td>
<td>1–10</td>
<td>0.5–10</td>
<td>0.1</td>
<td>1 [0.02]</td>
</tr>
<tr>
<td>density (cm⁻³)</td>
<td>&gt; 200</td>
<td>&gt; 10⁴</td>
<td>10³–10⁴</td>
<td>10³–10⁵</td>
<td>10⁶–10⁷</td>
</tr>
<tr>
<td>mass (M⊙)</td>
<td>10⁴–10⁶</td>
<td>100–10⁴</td>
<td>50–1000</td>
<td>0.5–5</td>
<td>10–1000</td>
</tr>
<tr>
<td>temperature (K)</td>
<td>10–30</td>
<td>&lt; 25</td>
<td>10–20</td>
<td>&lt; 15</td>
<td>&lt; 25</td>
</tr>
</tbody>
</table>

aRanging from small dark clouds to Giant Molecular Clouds
bThe size in brackets, as well as the density represent values for individual pre-stellar cores in a cluster.

temperature is \( \sim 10–30 \) K and the mean number density is in excess of \( \sim 200 \) cm⁻³, but, because of the hierarchical build-up these properties may vary significantly between and throughout the clouds.

Dark clouds are nearby molecular clouds that are observed in extinction at optical wavelengths (e.g., Ophiuchus). Dark clouds belong to the group of small molecular clouds described above. They are generally associated with sites of low- and intermediate-mass star formation, whereas the GMCs are often referred to as their massive, cluster-forming counterparts.

Infrared Dark Clouds (IRDCs) were discovered in the mid-90s independently by the Midcourse Space Experiment (MSX; Egan et al. 1998) and the Infrared Space Observatory Camera (ISOCAM; Pérault et al. 1996). They appear as dark, often filamentary structures against a bright mid-infrared background. IRDCs are presumably the direct precursors of clustered, massive star formation and they play an important role in this thesis. Their properties are more extensively discussed in Chapter 2.

Clumps

On scales of \( \sim 0.5–10 \) pc molecular clouds often contain filamentary, over-dense substructures referred to as clumps. Their masses range from \( \sim 50 \) to 1000 M⊙, the gas temperatures are \( \sim 10–20 \) K and the number densities range from \( \sim 10³ \) to \( \sim 10⁴ \) cm⁻³.

Cores

Pre-stellar cores are small, gravitationally unstable clouds where eventually a single star or stellar system is formed. Core masses range between 0.5 and 5 M⊙. Their size is of the order of 0.1 pc with a mean number density of \( \sim 10⁴–10⁵ \) cm⁻³. The temperature in the cores can be as low as about 8 –12 K.

Pre-cluster cores are the massive counterparts of pre-stellar cores. With masses ranging from \( \sim 10 \) to 1000 M⊙ (or more) they represent the precursors of small, loosely bound stellar clusters as well as massive OB-associations. Pre-cluster cores typically show a lot of substructure and the individual fragments appear to be precursors of single stars or stellar systems. The pre-stellar cores in cluster-forming regions
are more compact (∼0.02 pc) and denser (∼10⁶–10⁷ cm⁻³) compared to isolated pre-stellar cores (Ward-Thompson et al. 2007, and references therein). Pre-cluster cores are sometimes also called star-forming clumps.

**Dark Globules**

Dark globules, also referred to as Bok globules, are isolated, gravitationally bound molecular clouds that may have been removed from their parent molecular cloud by external events. Because of their isolation, the distance and consequently distance-related properties of many globules are often uncertain. Estimates of their mean size do not exceed ∼1 pc (Evans 1999, and references therein). Their mass is of the order of 10–100 M☉. The densest parts of globules resemble the pre-stellar cores described above. Low-mass star formation is frequently observed inside.

### 1.1.2 Relevant time-scales

The equations presented in this section give an idea of the time-scales involved in the evolution from molecular clouds to stars, though they are not meant to give the complete picture. They are expressed in units that are conveniently scalable to values appropriate for various objects. It is important to realise that mass (and density) play a crucial role in the time span of pre-stellar as well as stellar evolution.

Density perturbations can move through molecular clouds at a velocity limited by the speed of sound. The isothermal sound speed is, for typical conditions in the ISM, approximately given by (e.g., Spitzer 1998)

\[ c_s = \left( \frac{kT}{\mu m_H} \right)^{\frac{1}{2}} \text{km s}^{-1}, \tag{1.1} \]

where \( k \) is the Boltzmann constant, \( T \) is the temperature of the gas, \( m_H \) is the mass of a hydrogen atom and \( \mu=2.33 \) is the mean molecular weight of a particle, consistent with a 25% mass fraction of helium. Typical values in molecular clouds range from about 0.1 to 10 km s⁻¹. The time required for a cloud to approach a dynamical equilibrium state is of the order of the sound crossing-time. This is the time required to traverse the extent of a cloud at the sound speed, i.e.,

\[ t_{\text{cross}} \approx 10^7 \left( \frac{D}{10 \text{ pc}} \right) \left( \frac{c_s}{1 \text{ km s}^{-1}} \right)^{-1} \text{years}, \tag{1.2} \]

where \( D \) is the diameter of a cloud.

In the presence of a magnetic field, additional hydro-magnetic perturbations will move transverse and along the direction of the field lines at the Alvén velocity, given by

\[ V_A = 1.3 \left( \frac{B}{1 \mu \text{G}} \right) \left( \frac{1 \text{ cm}^{-3}}{n} \right)^{\frac{1}{2}} \text{km s}^{-1}, \tag{1.3} \]

where \( n \) is the number density. The field strength \( B \) ranges from about 0.001 \( \mu \text{G} \) in the diffuse ISM to \( \gtrsim 10 \mu \text{G} \) in dense cores. Observations show that \( B \) scales approximately with the density squared in molecular clouds. Therefore, a typical Alvén
velocity is of the order of 1 km s\(^{-1}\). In dense and cold cores, the Alvén speed can exceed the sound speed and become the limiting velocity for perturbations to traverse the medium.

The physical state of molecular clouds is heavily debated amongst experts in the field. It is beyond the boundaries of this thesis to go into the details. I will briefly review the two main directions here.

– quasi-static view –
The first is, that molecular clouds are quasi-static objects formed from the atomic gas phase due to gravitational instabilities. The quasi-equilibrium state of a cloud is maintained by the magnetic field through a magnetic force counteracting gravity and through support by hydro-magnetic waves. A cloud still evolves due to gravity because only charged particles feel the presence of the magnetic field. The majority of the gas is neutral (ionisation fraction \(\chi_i \sim 10^{-7}\)) and will diffuse inward, slowed down by collisions with the charged particles. This process is known as ambipolar diffusion and occurs on a time-scale given by

\[
t_{AD} = \frac{3}{4\pi G \rho \tau_{ni}} \approx 7.3 \times 10^6 \left(\frac{\chi_i}{10^{-7}}\right) \text{ years},
\]

where \(\tau_{ni}\) is the ion-neutral collision time, \(\rho\) the mass density and \(G\) the gravitational constant.

– dynamic view –
The opposite view is that molecular clouds are transient objects. They form from convergent flows driven by supersonic turbulence and evolve due to an interplay between gravity and turbulent motions. Hence this process is often referred to as gravo-turbulence. Turbulence dissipates on short time-scales. For supersonic turbulence this is approximately given by (Mac Low & Klessen 2004)

\[
t_{turb} \approx \frac{L_d}{v_{rms}} = 10^6 \left(\frac{L_d}{10 \text{ pc}}\right) \left(\frac{v_{rms}}{10 \text{ km s}^{-1}}\right)^{-1} \text{ years},
\]

where \(v_{rms}\) is the root-mean-square velocity and \(L_d\) the driving scale of the perturbations. The turbulent dissipation-time is short compared to the crossing-time and suggests that molecular clouds never reach a dynamical equilibrium. In the gravo-turbulent view, clouds are dynamic, transient objects.

For more information on the physical state of molecular clouds I suggest the interested reader to look through McKee & Ostriker (2007) and Bergin & Tafalla (2007) for a concise review and, moreover, follow up on the numerous references they give to original papers on the subject.

The onset of star formation
Initially, a combination of rotation and thermal, magnetic and turbulent pressure in a dense core withstands the inward gravitational force. Once the magnetic and turbulent support dissipate (on timescales given by Eq. 1.4 and Eq. 1.5) gravity overtakes the thermal pressure. A runaway process starts and a spherical cloud will
gravitationally collapse on a time-scale characterised by the free-fall time given by (Shu et al. 1987)

\[
t_{ff} = \left( \frac{3\pi}{32G\rho} \right)^{\frac{1}{2}} \simeq 2.4 \times 10^6 \left( \frac{n}{200 \text{ cm}^{-3}} \right)^{-\frac{1}{2}} \text{ years}.
\]  

(1.6)

An exact derivation of the collapse criterion, considering only thermal imbalance, can be derived using the virial theorem, defined as \(2 \times E_{\text{kinetic}} + E_{\text{gravitational}} = 0\). Collapse occurs when the gravitational energy exceeds twice the kinetic energy, with

\[
E_{\text{kinetic}} = \frac{3}{2} n k T \quad \text{and} \quad E_{\text{gravitational}} = \frac{-3GM^2_{\text{cloud}}}{5R_{\text{cloud}}}.
\]

(1.7)

where \(M_{\text{cloud}}\) is the cloud mass and \(R_{\text{cloud}}\) is the cloud radius. This condition can be approximated using the simple argument that the free-fall time should be less than the sound-crossing time. The critical value is usually expressed in terms of mass, referred to as the (thermal) Jeans mass (Jeans 1928). The collapse starts when the mass of a cloud exceeds the Jeans mass, i.e., \(M_{\text{cloud}} > M_J\), where the critical mass is approximately given by;

\[
M_J = \left( \frac{\pi c_s^2}{G} \right)^{\frac{3}{2}} \rho^{-\frac{1}{2}} \simeq 7.5 \left( \frac{T}{10^4 \text{ K}} \right)^{\frac{3}{2}} \left( \frac{n_{\text{H}_2}}{10^4 \text{ cm}^{-3}} \right)^{-\frac{1}{2}} M_\odot.
\]

(1.8)

The definition of the Jeans mass can be extended (see e.g., Larson 1998; Elmegreen 1999) by including turbulent support,

\[
M_J \propto v_{\text{rms}}^4 P^{-1/2},
\]

(1.9)

where \(P\) is the pressure; or magnetic support,

\[
M_J \propto B^3 \rho^{-1}.
\]

(1.10)

While the gas accumulates at the centre of a cloud, the core continues to contract. Gravitational potential energy is converted to heat, increasing the core temperature until hydrogen fusion can begin. The time-scale in which the pre-hydrogen-burning evolution occurs is given by the Kelvin-Helmholtz time, i.e.,

\[
t_{KH} = \frac{GM^2}{R_* L_*} \simeq 3 \times 10^7 \left( \frac{M_*}{1M_\odot} \right)^2 \left( \frac{R_*}{1R_\odot} \right)^{-1} \left( \frac{L_*}{1L_\odot} \right)^{-1} \text{ years},
\]

(1.11)

where \(M_*\) is the stellar mass, \(R_*\) the stellar radius and \(L_*\) the stellar luminosity. For a solar type star this stage of the evolution is relatively long, about a few times \(10^7\) years. For more massive stars however (\(> 8 M_\odot\)), the free-fall time on which the collapse occurs is longer than the Kelvin-Helmholtz time and accretion may continue while the star has already reached the main sequence phase.
1.1.3 The formation of isolated, low-mass stars

Figure 1.1: This figure illustrates the different evolutionary stages during the formation of an isolated, low-mass star, as described in the text. It is adapted from van Dishoeck & Blake (1998).
Theoretically, the formation-process of low-mass isolated stars ($\sim 1 M_\odot$) is well understood (Shu et al. 1987). To a certain extent the scenario is observationally confirmed by numerous studies in nearby low-mass star-forming regions. Starting initially as a dense and cold pre-stellar core, the evolutionary sequence is subsequently described by four main stages (Lada 1987). The shape of the near- to far-infrared spectral energy distribution (SED), dominated by the dust emission, is used to classify the stages as class 0, I, II or III objects (e.g., Lada & Wilking 1984; Andre & Montmerle 1994). Figure 1.1 illustrates the stages which are briefly described hereafter. The characteristic shapes of the SEDs are schematically drawn in Figure 1.2.

**Figure 1.2:** Illustration of the evolution of the SED during pre- and proto-stellar stages. Adapted from (Lada & Wilking 1984; Andre & Montmerle 1994).

- **Class 0** – The SED resembles a blackbody spectrum at a temperature below 30 K, peaking at sub-mm wavelengths ($\gtrsim 100 \mu m$).
- **Class I** – The peak of the SED shifts to far-IR wavelengths ($\lesssim 100 \mu m$) as the temperature of the dust rises. Emission from both the accreting envelope ($\sim 100 K$) and the thick disk (a few 100 K) are observed. The 10 $\mu m$ silicate absorption feature is often identified toward Class I objects, indicating that the envelope is optically thick.
- **Class II** – The SED at near-IR wavelengths is dominated by the emission of the central proto-star and typically peaks around 2 $\mu m$, corresponding to temperatures around 1000 to 2000 K. At longer wavelengths an infrared excess is observed, originating from the dusty disk.
- **Class III** – The SED resembles a stellar blackbody, peaking at optical or infrared wavelengths. Minor infrared excess is often still observed. Possibly, this is due to the remaining thin disk, but it is also consistent with interstellar extinction due to foreground material.
Dark cloud – In a typical dark cloud, several pre-stellar cores with enhanced densities are present. In the scenario of isolated star formation, these sub-structures are spatially well separated so that they do not interact with each other. Initially, rotation and internal pressure prevent a pre-stellar core from collapse under its own gravity. The core evolves slowly because of ambipolar diffusion (Eq. 1.4) and dissipation of turbulent motions (Eq. 1.5) due to collisions between particles (viscous effects). On time-scales of a few million years the core develops a strong increase in density toward the centre.

Class 0 – Defined at time $t=0$, the central density of the core becomes so high that gravity overcomes any internal pressure and the core starts to collapse. Due to the collapsing gas, a shock wave propagates outward at the speed of sound and the surrounding material starts falling in as well. This is known as an inside-out collapse (Shu et al. 1987). The slow rotational motion of the core that is present from the start is enhanced as accretion proceeds. Observational evidence of bipolar outflows at these stages (e.g., Lada 1999) suggests a non-spherical collapse. This might indicate that the formation of a disk starts early in the formation-process.

Class I – After about $10^4$ to $10^5$ years, the collapsing core becomes so dense that cooling radiation is trapped and heats the environment. The temperature of the central object increases gradually to a few hundred degrees and a proto-star is formed. Due to rotation, the angular momentum forces the collapsing envelope to accrete onto a thick circumstellar disk. The material in the disk continues to spiral in and onto the proto-star, allowing it to grow in size and mass. Excess angular momentum is expelled into the surrounding medium in the form of collimated jets and outflows along the rotation axis.

Class II – About a million years after the initial collapse, most of the envelope has been removed and the embedded object becomes visible at optical and infrared wavelengths. At this stage, the bulk of the material has accreted onto the central object. A flattened circumstellar disk is present in which material moves inward at a decreasing rate. The disk contributes only about 1% of the total mass of the system. Planet formation may have started. Material from a remaining envelope may still accrete onto the outer parts of the disk. The object is referred to as a classical T-Tauri star, after the prototype T-Tauri. The luminosity of the star is driven by compression due to its own gravity. This stage initiates the pre-main sequence stage of a star.

Class III – At some point, between 1 to 10 million years, accretion has come to a stop completely. Most of the primordial material in the disk is lost. What remains is a debris disk. The temperature and density of the pre-main sequence star keep increasing as the object slowly contracts to its final size. Class III objects are sometimes called weak T-Tauri stars.

Main sequence star – When the temperature in the centre of the star reaches about 10 million degrees, hydrogen fusion begins. The star settles onto the main sequence evolutionary track and will continue burning hydrogen for several billion years.
1.1.4 The formation of massive stars and stellar clusters

The formation of massive stars is poorly understood compared to their low-mass equivalents. From the observational point of view, massive star formation is more difficult to study in detail for a number of reasons. First of all, massive stars are rare (see also Section 1.1.5). Therefore, their formation sites are located at large distances ($\gtrsim 1 \text{kpc}$) making it difficult to spatially resolve them. Second, the lifetime of massive stars, including the formation-process, is shorter by orders of magnitude ($\sim 10^5$ years, see Section 1.1.2). To make things even more complicated, the formation-process occurs deeply embedded in the densest parts of molecular clumps, hidden almost completely from view.

Theoretically, the difficulty is that massive stars form almost exclusively in stellar clusters (Blaauw 1964; de Wit et al. 2005). Complications in the clustered star formation process are among other things the role of feedback from massive stars on their environment and the strong interaction between different star-forming events due to their proximity.

Young massive star clusters in our Galaxy have a stellar mass-content of the order of $1000 M_\odot$ (e.g., Orion OB1) to $10,000 M_\odot$ (e.g., Arches cluster). Clearly, such massive clusters must originate from molecular clouds that have sufficient material available to transform large amounts of mass into stars, the Giant Molecular Clouds. Inside these massive clouds, dense filaments fragment and collapse to form dense pre-cluster cores. On spatial scales of the order of the size of pre-cluster cores ($\lesssim 0.1 \text{pc}$), a number of subsequent stages in the formation-process of clusters have been identified, i.e., cold dense clumps, hot cores, hyper- and ultra-compact HII regions and OB-associations.

On even smaller spatial scales, the individual, dense, collapsing cores produce stars of all masses that will give rise to the stellar cluster. The exact formation-process of individual stars in a cluster, specifically the massive ones, is strongly debated. Various scenarios have been suggested. I will review three popular theories that have been proposed recently:

- **Competitive Bondi-Hoyle accretion** (Bonnell et al. 1997, 2001)
  In dense clusters, accreting objects are competing for the same material that is available. Simulations show that two different regimes are recognised in the accretion process. One is, when the gravitational potential is dominated by cold gas. In this regime, low-mass stars accrete the majority of their mass. The other is when the stars that have formed dominate the gravitational potential. At this stage, high-mass stars accrete most of their mass. The latter occurs first at the centre of a cluster, which explains why the most massive stars are usually found in cluster cores (Bonnell & Davies 1998). This model accounts for the observed mass distribution of stars more massive than $0.5 M_\odot$. However, for stars more massive than $10 M_\odot$, the radiation pressure of the central object suppresses further accretion.
• **Stellar mergers** (Bonnell et al. 1998)
  The merging mechanism is proposed to avoid the problem of direct accretion of material onto proto-stars more massive than about $10 \, M_\odot$, where radiation-pressure becomes too large. Instead, in the central parts of a cluster, intermediate-mass stars form through accretion onto lower mass objects as the cluster itself decreases in radius due to global collapse. At some point, the central density becomes so large that the intermediate-mass objects merge and form massive stars. This process can occur only at very high stellar densities ($\gtrsim 10^8 \, \text{pc}^{-3}$), which have not been observed so far. The model also predicts a large number of binary stars due to tidal captures. Note here, that massive stars indeed have a high degree of multiplicity compared to low-mass stars, i.e., an average massive stars has 1.5 companions against 0.5 for a solar-type star.

• **Turbulent core model** (McKee & Tan 2003)
  This model extends the low-mass formation picture that stars form through accretion onto a central object. This model suggests that massive molecular cores are supersonically turbulent even on the small scales where individual stars form. This implies an accretion rate that is high enough to overcome any radiation pressure of the central object. Furthermore, the accretion rate increases with time and can account for the short formation time-scale ($\sim 10^5$ years) of massive stars.

Up to now, the issue of how massive stars are formed remains unresolved. Apparently, different evolutionary stages of individual stars are present at the same instant in a cluster environment. The proposed evolutionary scenario of massive, clustered star formation is schematically given in Figure 1.3. The description of the different stages presented below is based on observational knowledge and is far from complete. It is meant to be merely descriptive.

**GMC** – Largely due to turbulent compression, driven by supernova shock waves, molecular clouds show a substructure that is highly filamentary. The gradual collapse under gravity and further compression due to shocks results in the formation of massive, dense clumps on scales of a few parsecs. These clumps are considered the pre-cursors of stellar clusters. Manifestations of cold clumps are presumably represented by the cold, centrally peaked sub-millimetre clumps observed in many IRDCs (see also Chapter 2).

**Collapse** – Defined at time $t=0$, gravitationally unstable pre-cluster cores of sizes less than $\sim 1 \, \text{pc}$ and masses ranging from 100 to $1000 \, M_\odot$ start to collapse rapidly. Initially, the temperature in the cores is still low ($\lesssim 20 \, \text{K}$). The clumps fragment further into pre-stellar cores that presumably undergo strong interaction with each other due to their small separation. A complicated, so far unresolved process (see the aforementioned formation mechanisms for individual massive stars) of clustered star formation starts to develop.
Figure 1.3: This figure displays a schematic illustration of the various observed stages during the formation of a massive stellar cluster, as described in the text.
**Hot core** – While the clump in whole is probably still collapsing, the excess angular momentum of individual proto-stellar objects drives jets and outflows into the surrounding envelope. Due to heating from the outflows, and from the proto-stellar cores, the temperature near the star forming activity in the clump rises to a few 100 K. This is sufficient to evaporate many of the ices from the dust grains. This hot region is about 0.1 pc in size and is observationally characterised by a rich gas-phase chemistry.

**HC/UC-HII** – Perhaps already during the hot core phase, the first massive proto-stars begin to fuse hydrogen. Their strong radiation ionises the gas and destroys complex molecules in their immediate surroundings. Observationally, this stage is identified as a hyper-compact HII region (HC-HII). The size of the ionised region is initially less than \( \sim 0.01 \) pc. As the ionisation front expands (< 0.1 pc) and further accretion ceases, the massive star reaches its final stage. The cluster is still embedded in an envelope of gas and dust and is referred to as an Ultra-Compact HII region (UC-HII).

**OB-stars** – About 150,000 years after the initial collapse, the HII region breaks through the surrounding envelope and the cluster of stars becomes visible. The presence of massive type O- and B-stars in the cluster gives rise to the name OB-association. The immediate surrounding of massive stars consists mainly of ionised gas and is prominently observed as classical HII regions.

### 1.1.5 The Initial Mass Function

Massive stars produce \( \sim 99\% \) of the UV photons in the Galaxy and are likely to be the dominant sources that ionise large parts of the ISM (the warm and hot ionised gas phase). However, massive stars are rare. By number they account for only \( \sim 1\% \) of all stars. The function that describes the distribution of stellar masses for a given stellar population is referred to as the Initial Mass Function (IMF). Observational evidence suggests that the IMF is very similar in most directions in our Galaxy. The IMF is defined as the number of stars \( \xi(m) \, dm \) that are produced in a mass interval \( m \) to \( m+dm \) and can be written as a power-law

\[
\xi(m) \propto m^{-\alpha},
\]

where \( \alpha \) is defined as the power-law value. Salpeter (1955) identified a power-law with \( \alpha = 2.35 \) for stars more massive than \( 1 \, M_\odot \). Hence, the high-mass end of the mass distribution is often referred to as the Salpeter function. Miller & Scalo (1979) showed for stars in the solar neighbourhood that the distribution breaks around \( 1 \, M_\odot \) objects, where the slope turns shallower. With the sensitive telescopes available at present, the low-mass end of the mass distribution can be studied as well. A standard IMF that is often used and covers the whole range of observed masses is given by Kroupa (2001) and consists of three parts (see Figure 1.4):

\[
\alpha = 2.3 \quad 0.5 < \frac{m}{M_\odot} < 50
\]
\[ \alpha = 1.3 \quad 0.08 < \frac{m}{M_\odot} < 0.5 \] (1.13)

\[ \alpha = 0.3 \quad 0.01 < \frac{m}{M_\odot} < 0.08 \]

There are indications that the IMF is not as universal as initially thought. In our own Galaxy, variations have been identified toward star clusters located near the Galactic centre (Scalo 1998). In addition, in external galaxies there is evidence that the stellar mass distribution deviates from the standard IMF as observed in our Galaxy. For example, a detailed study of young super star-clusters in the nuclear starburst galaxy M82 (McCrady et al. 2005) reveals variations of the IMF, although mass segregation is an issue.

1.2 Observing the early stages of star formation

From the previous sections, it is clear that the initial stages of clustered star formation are represented by the cold, dense and massive clumps inside Giant Molecular Clouds. On many occasions, distinct regions in the same GMC represent different evolutionary stages of star formation. The main subject of this thesis is an investigation of the early stages of clustered star formation, however, more evolved regions are often encountered in the observations.

In this Section I briefly discuss the observational possibilities that enable a detailed analysis of the physical conditions in different stages of star formation. I will focus mainly on observations that have been essential for the motivation of this thesis.

1.2.1 Interstellar dust

The mass of a molecular cloud consists for about 1% of dust. In general, silicates and graphite are considered the main constituents of interstellar dust. The dust grains have a broad range in size, from only a few molecules sticking together to micrometer sized aggregates. A significant fraction of the total mass of dust is contributed by large grains, i.e., \( \gtrsim 0.01 \mu m \).
Emission characteristics

Except for the smallest grains, interstellar dust is in general close to being in thermal equilibrium with its surroundings. The temperature of dust grains is regulated by various heating mechanisms, depending on the location in a cloud. The interstellar radiation field (ISRF) is the dominant heating source in the outer, lower density layers of clouds. The grains absorb the stellar radiation and re-emit their energy at typical temperatures around 30 K. Going deeper inside the cloud, most of the ISRF-photons are absorbed and the temperature drops. At high densities (>10^4 particles cm^{-3}), the dust is thermally coupled to the gas through collisions. The heating is only due to cosmic rays that penetrate the dense cores and ionise the gas. Even though the energy input per cosmic ray can be enormous, the low rate at which events occur (∼10^{-17} s^{-1}) yields a temperature of only 8 to 10 K. On small spatial scales, temperatures can increase due to local events such as compression, shocks, outflows and radiation from nearby (proto-) stars.

Large grains (≳ 0.01 μm) can be treated as blackbodies. The re-emitted radiation of large grains is observed as a modified blackbody spectrum. The intensity I as a function of wavelength λ is expressed as the Planck law multiplied by M(λ);

\[ I_\lambda(T) = M(\lambda) \frac{2hc^2/\lambda^5}{\exp(hc/\lambda kT) - 1}. \] (1.14)

Here h is the Planck constant, c is the speed of light, k is the Boltzman constant and T the temperature of the dust. The modification factor can be written as

\[ M(\lambda) = 1 - \exp \left[ -\tau_\lambda(0) \left( \frac{\lambda(0)}{\lambda} \right)^\beta \right], \] (1.15)

where \( \tau_\lambda(0) \) is the optical depth at a reference wavelength \( \lambda(0) \) and \( \beta \) is a correction factor depending on the intrinsic dust properties, mainly the size-distribution and composition. Values for \( \beta \) may vary between about unity and 3. Considering the low temperatures in molecular clouds, the majority of the energy is emitted at far-infrared and sub-millimetre wavelengths. The radiation at these long wavelengths is optically thin (τ ≪ 1, i.e., radiation can freely escape the cloud). Therefore, the observed intensity is a direct measure of the total column of radiating dust along the line of sight. The wavelength at which the peak of emission occurs (\( \lambda_{\text{max}} \)) relates to the temperature (\( T_{\text{dust}} \)) of the dust and is given by Wien’s Law;

\[ \lambda_{\text{max}} T_{\text{dust}} = 0.29 \text{ cm K} \] (1.16)

Small dust grains (i.e., ∼20–50 Å) cannot be considered as blackbodies. They readjust to higher temperatures compared to large grains and emit their energy at infrared wavelengths, typically around ∼60 μm for small grains and ∼25 μm for the very small grains. The Polycyclic Aromatic Hydrocarbons (PAHs) account for most of the grains at molecular size scales (4–20 Å). These grains are not in thermal equilibrium with their surroundings. A random absorption of photons heats the individual grains to a range of temperatures. The energy is re-emitted at very specific
wavelengths, predominantly in the near- to mid-infrared, depending on the internal properties of the grains (e.g., Tielens et al. 1999). Frequently used models that describe the grain size distribution and dust composition of the ISM as well as present synthetic dust emission spectra are, e.g., Draine & Li (2001, 2007); Weingartner & Draine (2001).

**Extinction characteristics**

The presence of dust in the ISM is prominently observed at optical and infrared wavelengths due to the absorbing and scattering properties of dust grains. Star counts were recognised as early as in the mid-20s of the previous century as a powerful tool to map nearby dark globules (Wolf 1923). The amount of extinction at wavelength $\lambda$, usually converted to magnitudes of visual extinction ($A_V$), can be derived directly by comparing the stellar distribution behind the dust with an extinction-free off-position, i.e.,

$$A_V = \frac{1}{b_\lambda} \frac{A_V}{A_\lambda} \log \frac{N_{\text{off}}}{N_{\text{on}}} \text{ mag},$$

(1.17)

where $b_\lambda$ is the slope of the cumulative $\lambda$-luminosity function, $N_{\text{off}}$ is the stellar surface density at the off-position and $N_{\text{on}}$ is the stellar surface density in the target direction. The ratio $A_V/A_\lambda$ is the reddening-law and depends on the dust properties. Star counts require a completeness-, or magnitude-limited, set of observations and assume that all observed stars are behind the obscuring material.

In particular at infrared wavelengths, the colour excess of objects behind the dust can be used to measure the extinction. The colour excess is defined as the difference between the observed colour and the intrinsic colour of a source, e.g., for $H - K_S$ the colour excess is given by

$$E(H - K_S) = (H - K_S)_{\text{observed}} - (H - K_S)_{\text{intrinsic}} \text{ mag.}$$

(1.18)

Often, the intrinsic colour of individual sources is unknown. However, the near-IR $H - K_S$ colour spans only a narrow range for the majority of stars, i.e., about 0.00 to 0.30 mag for spectral types A0V to M6III (Bessell & Brett 1988) with a mean of $\sim 0.15$ mag. Typical $H - K_S$ extinction values toward the denser parts of molecular clouds are in excess of 0.3 mag, i.e., a mean $(H - K_S)_{\text{observed}} > 0.45$ mag. The infrared colour excess is often converted to visual extinction by multiplication with a constant. This conversion assumes that the reddening-law is known.

**1.2.2 Molecular gas**

A continuation of the initial collapse of a molecular cloud requires that the material cools down to suppress the thermal and turbulent support. Molecules are very effective in cooling the gas, because of the presence of many rotational energy levels. Transitions between different levels involve only small amounts of energy, i.e., for many molecules the lowest transitions require less than $\sim 10$ K. The emitted radiation is therefore dominant in the (sub-) millimetre regime, unlike vibrational or electronic transitions which emit predominantly at infrared and optical wavelengths, respectively. Because of these low energies, the first rotational levels of molecules are
easily populated, even in cold regions. Most of the spectral lines observed toward cold molecular gas are due to pure rotational transitions.

**Rotational transitions**

In this thesis mainly the rotational transitions of diatomic molecules (mostly CO) are used. The energy levels are characterised by the rotational quantum number \( J \), which denotes the magnitude of the total angular momentum. For diatomic (and linear) molecules, the energies are approximately given by the levels allowed for a rigid rotator, i.e.,

\[
E_J = B J (J + 1), \quad J = 0, 1, 2, 3...
\]

(1.19)

The rotational constant \( B \) is given by

\[
B = \frac{\hbar^2}{8\pi^2 I},
\]

(1.20)

where \( \hbar \) is the Planck constant and \( I \) the moment of inertia of the molecule. An observer detects the photons that are emitted due to radiative de-excitation from an upper to a lower level. The notation that is generally used, for example for the transition between the \( J = 2 \) and \( J = 1 \) level of carbon monoxide is CO 2-1. For radiative rotational transitions, the only allowed transitions between different \( J \) levels obey the selection rule

\[
\Delta J = \pm 1.
\]

(1.21)

The energy corresponding to a transition is given by

\[
E(J + 1) - E(J) = 2B(J + 1).
\]

(1.22)

From this equation, it follows that the rotational spectrum of a molecule consists of evenly spaced lines in the frequency domain. Also, massive molecules have a large inertial momentum \( I \) and thus, the energy levels are more closely spaced than less massive molecules.

**Critical density, optical depth and intensity**

The energy level populations are determined by solving the statistical equilibrium between the excitation and de-excitation of the molecules due to both radiation and collision. Radiation can excite molecules through absorption, and de-excite molecules through spontaneous or stimulated emission. Collisional energy-exchange between molecules (or other collisional partners such as atoms or electrons) can also result in both excitation and de-excitation. In molecular clouds it is usually sufficient to take into account only molecular hydrogen as a collision partner.

In order to simplify the problem, consider a two-level molecule where subscripts \( u \) and \( l \) denote the upper and lower level, respectively. The equilibrium equation is then given by

\[
n_l n_u \gamma_{lu} + n_l B_{lu} J_{ul} = n_l n_u \gamma_{ul} + n_u A_{ul} + n_l B_{ul} J_{ul},
\]

(1.23)

where \( B_{lu}, B_{ul} \) and \( A_{ul} \) are the Einstein coefficients giving the absorption rate, the stimulated emission rate and the spontaneous emission rate, respectively. The \( \gamma \)'s
are the collisional rate coefficients. The number density is given by \( n \) and the mean photon radiation field is \( J_{ul} \). Solving this equation is not straightforward, because at any location in a cloud the level populations depend on the radiation field while the radiation field itself is determined (partly) by the level populations everywhere.

For a detailed evaluation of the equilibrium equation as well as the basics of radiative transfer I encourage the reader to have a look at Rybicki & Lightman (1986). I will only briefly touch upon some key-topics here to clarify two parameters that will be discussed further on in the thesis, i.e., the critical density and the optical depth.

An important parameter that is used when solving the equilibrium equation is the critical density \( n_{cr} \), i.e., the density at which the radiative decay rate equals the collisional de-excitation rate. If we ignore photon-trapping for the moment (optically thin limit), the critical density is defined as the ratio of the de-excitation rates through spontaneous emission and collision, respectively;

\[
\frac{n_{cr}}{\gamma_{ul}} = \frac{A_{ul}}{\nu^3}.
\]  

(1.24)

The Einstein \( A_{ul} \) coefficient is the radiative de-excitation rate. \( A_{ul} \) scales with the frequency cubed \( (A_{ul} \sim \nu^3) \) and thus, higher energy transitions have a higher critical density. The collisional de-excitation rate coefficient \( \gamma_{ul} \) depends on the de-excitation cross-section and the density of the collision partner. A useful implication of the critical density for observers is, that the detection of specific molecules and transitions can be used as constraint on the number density in the region they are observed. In particular, in Chapter 5 we make use of the detection of the C\(^{34}\)S 2-1 transition as a probe for high density \( (n_{cr} \approx 5 \times 10^5 \text{cm}^{-3}) \) cores.

In the case of an optically thick transition, some of the emitted photons are re-absorbed. The absorption is characterised by the absorption coefficient \( \alpha_{\nu} \). Re-absorption has no net effect on the level populations in a cloud, because a photon that is emitted due to de-excitation of one molecule excites another molecule elsewhere in the cloud. As a result, the critical density is lowered by \( 1/\tau_{\nu} \). Here, \( \tau_{\nu} \) is the optical depth of the line, defined as the integral of the absorption coefficient over the traversed path \( s \), i.e.,

\[
\tau_{\nu}(s) = \int_0^s \alpha_{\nu}(s')ds'.
\]  

(1.25)

At densities much below the critical density \( (n \ll n_{cr}) \), molecules are radiatively de-excited at a much faster rate than they are de-excited by collisions. At high densities \( (n > n_{cr}) \), collisions dominate the de-excitation process and the gas is in local thermodynamic equilibrium (LTE).

It is instructive to mention here the dependence of the emergent intensity in a line, i.e., the intensity that is observed by a telescope. In the optically thin limit \( (\tau \ll 1) \), the intensity scales linear with the column density in the upper level and thus can be used to estimate the total column of material along the line of sight. In case of an optically thick line \( (\tau > 1) \), the intensity is given by multiplying the Planck function
with the width of the line. This means the intensity is a direct measure of the excitation temperature $T_{\text{ex}}$ of the gas. This immediately gives an approximation of the kinetic temperature $T_{\text{kin}}$, because at high density $T_{\text{ex}} \simeq T_{\text{kin}}$.

**Line width – size relation**

The observed broadening of spectral lines toward molecular clouds is dominated on most spatial scales by supersonic motions. A power-law relation between the velocity dispersion $\sigma$ (km s$^{-1}$) inside a dark cloud and the physical extent $L$ (pc) was recognised from observations by Larson (1981) and is written as

$$\sigma = 1.1L^{0.38}.$$  \hfill (1.26)

The equation is empirically derived for molecular cloud sizes ranging from about 0.1 to 100 pc. The result is almost identical to what is expected on the basis of the classical Kolmogorov law of incompressible turbulence. Because no systematic evidence for other mechanisms is present, such as rotation or in- and outflow motions, the supersonic motions are generally ascribed to the presence of turbulence in molecular clouds.

**Molecular hydrogen or carbon monoxide**

Molecular hydrogen ($\text{H}_2$) is by far the most abundant molecule in the ISM. However, $\text{H}_2$ is a homonuclear molecule and does not have a permanent dipole moment. This means that it has no allowed dipole transitions. Weak quadrupole transitions ($\Delta J = \pm 2$) do exist, but the energies involved are high, e.g., $E_{J=2} - E_{J=0} \sim 510$ K. Consequently, $\text{H}_2$ is difficult to observe in molecular clouds and cannot be used to trace the molecular content of the Galaxy.

Carbon monoxide ($\text{CO}$) is the most abundant molecule in the ISM after $\text{H}_2$. It is the dominant coolant in cold gas and is widely used as a tracer of molecular gas. The lower rotational transitions require low energies (e.g., $E_{J=1} - E_{J=0} = 5.5$ K) to jump to the first excited rotational states. CO is observed everywhere in molecular clouds. Even though the abundance of CO is low ($\sim 10^{-4}$), the immense column densities in molecular clouds make CO molecules optically thick to their own radiation. This means that the observed CO emission is predominantly coming from the outer layers of molecular clouds. To acquire information on the entire column of CO along a line of sight, which may be converted to a mass column of gas assuming a standard abundance, it is essential to observe the rare isotopes, e.g., $^{13}\text{CO}$ and $^{18}\text{O}$ (abundance ratios are $\approx 50$ for $^{12}\text{C}^{13}\text{O}$ and $\approx 500$ for $^{16}\text{O}^{18}\text{O}$).

### 1.3 Thesis outline

Chapter 1 presents a general review of the current knowledge on star formation and introduces some of the basic physics related to the subject.

Chapter 2 gives an overview of the global properties of IRDCs. I describe the initial identification and spatial distribution toward the Inner Galaxy and give typical
values for the mass, density and temperature. At the end of the chapter I address the main topic of this thesis.

Chapter 3 introduces a statistical approach to identify extended red regions from the Two Micron All Sky Survey (2MASS) in the Outer Galaxy. I explain how we use the Mann-Whitney U-test and a cluster finding algorithm to produce a list of candidate dark clouds. The catalogue is cross-correlated with existing infrared and optical data.

Chapter 4 continues the identification of Outer Galaxy dark clouds by matching candidates with CO data from the FCRAO survey. The presence of CO suggests that the identified objects are red because the 2MASS sources are embedded in, or behind a molecular cloud.

Chapter 5 provides a detailed analysis of molecular line observations toward the dark cloud candidate G111.80+0.58 in the vicinity of NGC 7538. The global properties that I present suggest that the cloud is in an early, but not pristine stage of star formation. The cloud appears to be very similar to Inner Galaxy IRDCs and I conclude that this is very likely the first Outer Galaxy IRDC-like object identified as such.

Chapter 6 presents mid-infrared Spitzer observations of G111.80+0.58. The conclusion is simple. G111.80+0.58 is observed in extinction at 8 micron and thus can be classified as the first real IRDC in the Outer Galaxy. More Spitzer data are expected in the near future that will allow a detailed investigation of the proto-stellar properties of several young stellar clusters associated with the dark cloud complex.

Chapter 7 summarises the main conclusions of the thesis. I address some of the future prospects arising from this thesis.