5 Star Formation

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Abstract: Stars are one of the most important constituents of the Universe, and understanding their formation is crucial to many areas of astrophysics. Stars form from dense molecular gas, and they tend not to form in isolation. Stars often form in binary and multiple systems, and these systems tend to form in clusters with 10^2–10^5 members. Stars also form with a wide range of masses, from substellar brown dwarfs with masses <0.1 $M_\odot$ to massive stars >100 $M_\odot$, and wherever stars form the distribution of their masses seems always to be the same. This chapter will review our current understanding of star formation from cold gas to young star clusters.

Keywords: Initial mass function, Molecular clouds, Stars, Star formation, Star clusters

1 Introduction

How stars form is one of the big questions in modern astrophysics. Stars are, in many ways, the fundamental (baryonic) components of the Universe. Most of the electromagnetic radiation detected by our telescopes comes directly or indirectly from stars, they are the basic visible components of galaxies, they are the crucibles in which heavy elements are produced, and they are the hosts of planetary systems and even, in at least one case, life.

Understanding star formation requires us to understand gravity, turbulence, magnetic fields, chemistry, thermodynamics, and radiative transfer processes, all acting together with extremely complex interactions and interdependencies. The study of star formation involves understanding physical processes that work on galactic kpc-scales, star cluster formation on pc-scales, and star (and planet) formation on au- and stellar-scales, with mass ranges from whole galaxies of $>10^{12}$ $M_\odot$ to Jupiter-sized planets of only 10^{-3} $M_\odot$.

The enormous range of scales which need to be probed to understand star formation make it extremely challenging. Only in recent years have computers become powerful enough to start dealing with the problem of star formation, informed by observations which are now available (often only from space) across the whole electromagnetic spectrum.

This chapter will not even attempt to cover much of the theory and observations of star formation (entire books are not enough). Rather, it will give a general overview of some of the more important processes and our understanding of them.

1.1 The Basic Model of Star Formation

The basic theory of star formation dates back to 1734 when Emanuel Swedenborg first proposed the Nebula Hypothesis: that the Sun and the Solar System formed from a rotating cloud of gas that collapses under gravity to form the Sun at the center and a disk around it from which the planets formed. This model was expanded later in the eighteenth century by Immanuel Kant and Pierre-Simon Laplace. Perhaps surprisingly, the basic idea of the nebula hypothesis is still the basis of our understanding of star formation. Current thinking is that star formation can be divided into several main stages.

Molecular cloud formation. A large (10^3–10^6 $M_\odot$) cloud of gas forms within the interstellar medium. The high column and volume densities allow the cloud to cool, and remain cool, and molecular hydrogen (and other molecules) to form.

Prestellar core formation. The molecular cloud fragments into self-gravitating condensations known as "clumps" and, on smaller scales, "cores." "Prestellar cores" of roughly a solar mass are the birth places of stars.

Embedded star formation. Prestellar cores collapse and form a protostar (actually, often a binary or multiple system) surrounded by a disk of gas near their centers. Initially these protostars are large (AU-size) hydrostatically supported objects deeply embedded in the gas of the core.

Pre-main sequence stars. Once most of the mass in the core is accreted onto the star(s), and young stars on the pre-main sequence are observed to be surrounded by massive disks. These disks may well be in the process of planet formation and fairly rapidly disappear. Eventually the star will reach the main sequence and become a "normal" main sequence star.

Star clusters. A molecular cloud usually fragments into many cores, forming a star cluster of between 10^2 (arguably even just ten stars can comprise a "cluster") and 10^6 stars in only a few pc^3 in a few Myr.[1] The extreme densities in star clusters result in frequent encounters between stars and multiple systems which can destroy young multiple systems and disrupt circumstellar (and planet forming) disks.

The end of star formation. Once one or more massive ($>10\,M_\odot$) stars form, the input of energy from their UV radiation fields (and associated HII regions), stellar winds, and eventual supernovae will expel any gas that has not already formed stars. This prevents more star formation and often results in the destruction of the star cluster.

These stages do not necessarily happen one after the other. Obviously stars cannot form before a molecular cloud has formed, but prestellar cores and pre-main sequence stars are often observed in the same cluster, while at the same time massive stars may have started to clear some regions of the cluster of gas so stopping star formation while star formation is still occurring in other regions (indeed, the removal of gas from one part of a cloud may well induce star formation in another).

1.2 What Is a Star?

Before proceeding, it is useful to define what is meant by a "star." Formally a star is defined as an object that will, is, or has produced energy through hydrogen fusion in its core. This is any object with a mass in excess of the hydrogen-burning limit of $\sim 0.08\,M_\odot$. For the purposes of this chapter, the "will" in this definition is crucial as many of the young objects that are discussed are far from reaching the main sequence and beginning core hydrogen burning.

A brown dwarf is an object too small to fuse hydrogen, but large enough to undergo a short-lived deuterium-burning phase (which produces very little energy as deuterium is very rare). Brown dwarfs span the mass range between $10^{-2}\,M_\odot$ (about ten Jupiter masses) to $\sim 0.08\,M_\odot$. Planets are therefore objects with masses $<10^{-2}\,M_\odot$, but this chapter will be generally unconcerned with planets and objects of even lower mass.

While this formal distinction between stars and brown dwarfs exists, throughout this chapter we will generally assume that brown dwarfs and stars are fundamentally the same type of object. Star formation should have no reason to draw a distinction between brown dwarfs and

[1] The difference between a massive star cluster and a small dwarf galaxy is that a star cluster forms all of its stars at roughly the same time with roughly the same metallicity (a "simple stellar population"), while a dwarf galaxy has many episodes of star formation over many Gyr.

stars. Only the most massive objects in young star forming regions have yet fused hydrogen to generate energy (even if they eventually will), and the physics of star formation should find nothing special about the hydrogen-burning limit. Therefore every statement about the formation of low-mass stars (especially very low-mass stars) should also apply to the formation of brown dwarfs.

The most massive stars that are forming now have masses of 150–300 $M_{\odot}$. It is difficult to estimate the masses of the most massive stars as young, massive stars are deeply embedded in gas. By the time the most massive stars are easily observable, after 1–2 Myr, they have evolved significantly and may have lost >50% of their initial mass.

1.3 Open Questions

There are a number of critical open questions in star formation:

The nature of molecular clouds. Stars tend to form in massive gas and dust clouds with sizes of 10s pc and masses up to 10^6 $M_{\odot}$. These clouds appear to contain high levels of turbulence which drives and controls star formation. What is the source of their turbulence and what is the lifetime of molecular clouds?

The origin of the initial mass function (IMF). Observations suggest that stars form with the same distribution of masses everywhere. There are very few brown dwarfs, and very few very high-mass stars. Most stars (90%) are M-dwarfs with masses in the range 0.1–0.5 $M_{\odot}$. Interestingly, and strangely, the IMF appears almost universal – observations of star forming regions show that their IMF always seems to be approximately the same. Why is it universal? And why is the typical mass always around 0.2 $M_{\odot}$?

The origin of multiple systems. Many (possibly most) stars do not form as single, isolated, objects, but in multiple systems with two or more members. The distribution of separations between members of multiple systems is extremely wide, with some companions almost touching, while others have orbits of millions of years. Why do stars form in multiple systems? And how do they produce such a wide range of separations between companions?

Is star formation universal? Do stars everywhere form in the same way? Star clusters range in mass from 10^2 to 10^6 $M_{\odot}$. Do more massive clusters just form more stars, or is there a fundamental difference in how they form stars? Do some regions form more or different binaries to other regions, or is it always the same?

There are many good reviews of, and introductions to, star formation, in particular: the *Protostars and Planets V* volume from 2007, and many articles from *Annual Reviews of Astronomy & Astrophysics* (especially recently McKee and Ostriker 2007, and Zinnecker and Yorke 2007), as well as the textbook *The Formation of Stars* by Stahler and Palla from 2005.

2 From Gas to Stars

Stars form from the gravitational fragmentation of cold gas. In the early Universe there were no stars; as the Universe cooled and the first dark matter halos formed, gas was able to collapse and cool in these halos. The formation of the first stars in a zero metallicity environment was probably very different to the formation of stars today, and they were probably very massive, short-lived stars (this is beyond our remit, but see Stacy et al. 2011 and references therein).

The introduction of metals to the interstellar medium (ISM) by the first generation of stars created dust and massively increased the efficiency of cooling, and later generations of stars were able to form in much the same way that stars currently form.

2.1 Star Forming Regions

Stars usually form in large complexes. In *Fig. 5-1* we see the Carina Nebula, a massive star forming complex. The image is about 20 pc across and illustrates almost every feature of star formation. One of the most obvious features is the complex and filamentary nature of the gas throughout the region. This is due to a combination of supersonic turbulence in the gas and the feedback of energy from the most massive stars. Stars are forming throughout the region including many very massive stars. η Carina is a ~100 $M_{\odot}$ star found at the far left as a bright blob just to the left of the Keyhole Nebula.[2] To the middle right is the star cluster Trumpler 14, while to the far left is the cluster Trumpler 16 (less easily visible but containing η Carinae) which are only a few Myr old. Along the bottom is a molecular ridge with very obvious dust pillars carved by the feedback from massive stars in Trumpler 14.

The Carina Nebula is a particularly extreme star forming region forming some 10^5 $M_{\odot}$ of stars in several clusters. Many stars also form in midrange clusters such as the Orion Nebula Cluster (the fuzzy blob in Orion's sword) with masses of 10^3 $M_{\odot}$, and many in very small clusters with only around 10^2 members such as the Taurus star forming region. As is described later, each of these types of star cluster contribute roughly the same mass of stars to the general field population of galaxies.

Fig. 5-1
An amazing HST ACS image of the Carina Nebula which is a massive star forming region. The image covers roughly 20 pc (Credit: NASA, ESA, N. Smith (University of California, Berkeley), and The Hubble Heritage Team (STScI/AURA), Credit for CTIO image: N. Smith (University of California, Berkeley) and NOAO/AURA/NSF)

[2] There is a lovely clickable version of this image available at http://heritage.stsci.edu/2007/16/supplemental.html describing many of the features.

2.2 Molecular Clouds

The most abundant element in the Universe is hydrogen. Hydrogen is most often found in one of three forms: ionized (HII), atomic (HI), and molecular (H_2). The interstellar medium (ISM) can be broadly divided into three phases: the hot ISM with temperatures $>10^4$ K which contains mainly HII, the warm ISM with temperatures between about 100 and 10^4 K containing HI, and the cold ISM at <100 K containing mainly H_2.

Stars form in the cold phase in clouds of gas and dust known as molecular clouds (MCs). Most molecular gas is found in Giant Molecular Clouds which typically have sizes of 10–100 pc, and masses of 10^3–10^6 $M_\odot$, but some molecular gas is found in smaller clouds known as Bok Globules. MCs have high column densities, shielding them from the interstellar radiation field and cosmic rays. This means that their temperatures can become very low (only 10s K) which means that they can form molecular hydrogen and many other molecular species.

2.2.1 Observations of MCs

Clouds of molecular hydrogen are unfortunately remarkably hard to observe. In the optical they are only observable as "dark clouds" as MCs can have tens of magnitudes of optical extinction meaning it is impossible to see their contents or anything behind them (which is how they were first discovered). Observations in the IR are able to see embedded (proto)stars within MCs, but not the MCs themselves.

Obviously it would be extremely useful to observe the structure and kinematics of the gas within MCs. But molecular hydrogen in MCs is most often unobservable: at the low (10s K) temperatures of MCs H_2 has no easily excited states. The only possible way of observing cold H_2 directly is to use UV absorption along the line of sight from a massive star, unfortunately there are very few of these lines of sight available, and almost never where you would like them to be. Therefore "tracers" must be used: other molecules or dust in the MC which are assumed to trace the underlying H_2 distribution.

For a far fuller description of the physics and techniques of MC observations see Evans (1999) and early chapters of Stahler and Palla (2005).

2.2.2 Molecular Tracers

Even though H_2 is not excited at low temperatures, many other molecular species are, and these are observable at radio (and mm) wavelengths. Molecular lines can be extremely useful as line ratios can provide information on the local (kinetic) temperature, and line widths provide information on bulk flows and nonthermal motions.

The emission of a particular line from a molecule depends on both the temperature and *volume* density (i.e., collision rate). However, the strength of the emission also depends on the column density of material.

Probably the most common tracer is CO which has a number of easily excitable lines at low temperature in the radio. CO is destroyed at low column densities by UV radiation and cosmic rays and so it can only exist in high (column) density environments such as MCs, and the CO column density is assumed to trace the H_2 column density, therefore allowing CO to be used to map H_2. Due to its relatively high abundance CO lines can saturate at high column densities.

In such situations isonomers of CO can be used (such as $C^{18}O$) which are less abundant and so do not saturate. However, CO freezes-out onto dust grains at very low temperatures (<10 K), forming an ice mantle around the dust grain.

Other molecules are used to trace regions of high volume density, for example, CS, HCN, NH_3, and many others. Different molecules (and lines) can be used to trace regions of different density. NH_3 is a particularly common tracer for densities of $\sim 10^3\ cm^{-3}$, while HCN traces volume densities of 10^6–$10^8\ cm^{-3}$. Line profiles can also be used to detect signatures of outflows and infall; however, the details of using and interpreting molecular tracers are quite complex and involved and the reader is referred to Evans (1999) for more information.

The power of using several molecular tracers in tandem is that the general column density structure of molecular clouds can be mapped (usually in CO), and regions of high volume density can be located using other tracers. In addition, the thermal and kinetic properties of the gas can be determined.

2.2.3 Dust as a Tracer

MCs contain significant quantities of dust (estimated to be roughly 1% of their mass). Unlike molecular tracers, dust does not emit line radiation, rather dust particles are large enough to emit a thermal continuum. The dust has the same kinetic temperature as the surrounding gas as collisions with gas particles are able to keep the dust and gas in thermal balance. At the typical temperatures of MCs of 10–100 K the peak thermal emission from dust lies in the sub-mm at wavelengths of hundreds to tens of microns.

Thus the total sub-mm flux in a region provides a measure of the column density of dust (which is converted into a gas column density using a gas-to-dust ratio that is usually taken to be 100-to-1), and the spectral energy distribution provides the gas (kinetic) temperature.

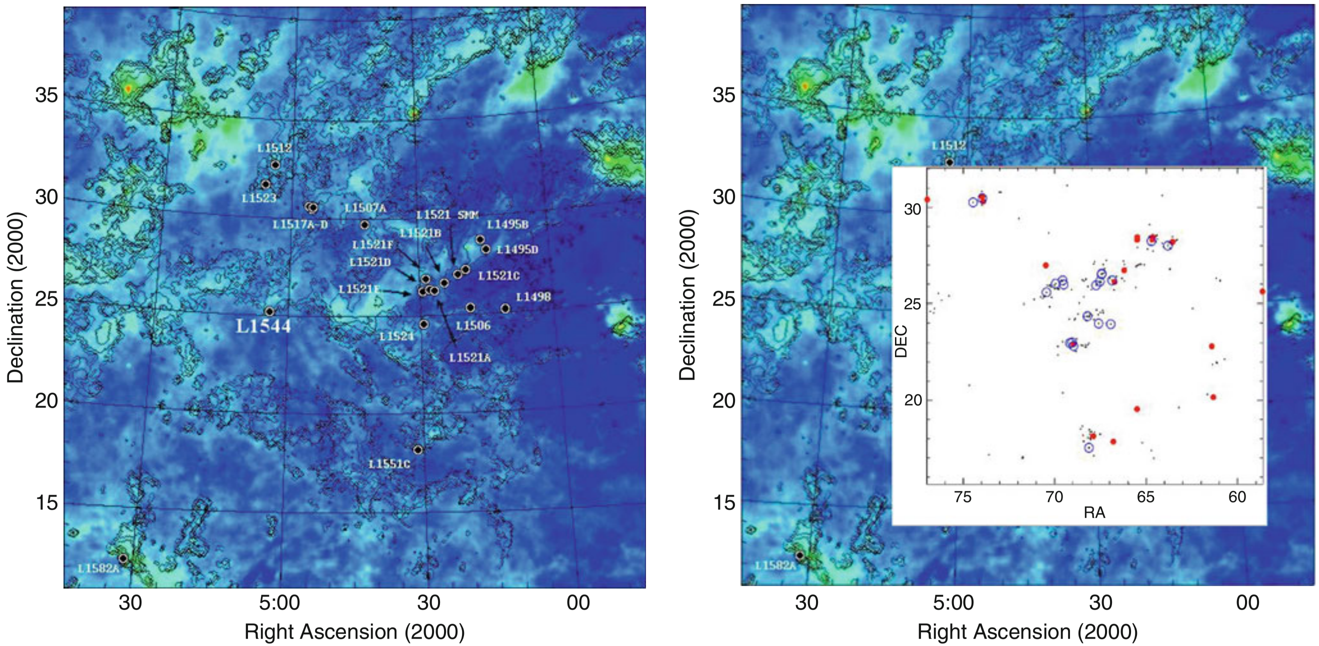
Unfortunately from the ground there are only a few atmospheric windows in the sub-mm, most notably at 450 and 850 μm, and so the whole sub-mm is not available (obviously this is not a problem from space and the Herschel mission is able to observe the whole of the sub-mm).

2.2.4 The Appearance of Star Forming Regions

In ➲ *Fig. 5-2* the gas and young stellar distribution in the Taurus star forming region is shown. The left panel from Kirk et al. (2005) shows the CO emission contours mapping the underlying H_2 distribution, while the color table shows the dust temperatures from dust continuum observations, from about 10 K for dark blue, to 200–300 K for yellow/red (e.g., at the top left). The black points show the positions of prestellar cores that are thought to be just about to form stars.

The right panel from Parker et al. (2011) shows the positions of the young stars in Taurus. Blue circles show the positions of brown dwarfs and very low-mass stars, red points the positions of stars >1 $M_\odot$, and the black points the positions of intermediate-mass stars.

There are several things to note about these figures which will become important in our discussions below. Firstly, the gas is not distributed evenly, it is clumpy and substructured with regions of very low density and filamentary structure in the gas. Secondly, prestellar cores – the sites of new star formation – are distributed in a similar way to the gas, generally forming



■ **Fig. 5-2**

***Left panel*: The gas distribution in Taurus from Kirk et al. (2005) showing CO emission contours, dust temperature in the background color table, and *black points* show the position of prestellar cores. *Right panel*: The stellar distribution in Taurus from Parker et al. (2011), *blue circles* show the positions of brown dwarfs and very low-mass stars, *red points* the positions of stars >1 $M_\odot$, and the *black* points the positions of intermediate-mass stars. Note that the projection of the two figures is slightly different**

where the gas is densest. Thirdly, young stars also follow the gas distribution, but less closely as they have had some chance to disperse from their formation sites, or for gas to clear from their formation sites.

That cores, and therefore stars, form where the gas is densest should not be a surprise. If the gas is substructured, then the initial stellar distributions in clusters will also be substructured. However, substructure does not remain for long. Most young clusters appear relatively smooth and circular which suggests that some process occurs very rapidly to remove the memory of the initial distribution of the stars/gas.

2.2.5 The Structure and Kinematics of MCs

Almost every MC is observed to have a large degree of internal density and kinematic structure down to the resolution limits of the observations. All MCs appear similar to Taurus (see *Fig. 5-2*) with clumps and filaments on all scales (Williams et al. 2000).

Often the structure in MCs is divided into "clouds," "clumps," and "cores":

Clouds are the largest structures representing the whole molecular cloud with typical sizes of pcs, masses of 10^3–10^6 $M_\odot$, and densities of ~100 cm^{-3}.

Clumps are denser regions within clouds which are associated with star cluster formation. They have sizes of around a pc, masses of a few hundred $M_\odot$, and typical densities of 10^3–10^4 cm^{-3}.

Cores are the sites of star formation in single or multiple systems with sizes of < 0.2 pc, masses of around a $M_\odot$, and densities of $>10^4$ cm^{-3}.

It should be noted that most of the mass in MCs is at relatively low density, clumps and cores only makeup a few percent of the total mass of a MC. The distinctions between clouds, clumps, and cores are rather arbitrary, and in many ways the structure in MCs appears scale-free which has led to suggestions that the structure of MCs is fractal (e.g., Elmegreen 2002).

This of course raises the question as to why MCs have such complex (fractal even?) structure? The answer is probably that MC structure is created and maintained by turbulence.

Observations of molecular line widths are a probe of the velocity dispersions of MCs. Line ratios show that the kinetic temperatures of MCs are fairly uniform, and very low, with typical temperatures throughout a MC of <100 K, down to only 10–20 K. However, line widths show significant supersonic motions in MCs. At 10–20 K, the sound speed of the gas is around 0.2 km s^{-1}; however velocity dispersions of many km s^{-1} are not unusual in MCs.

Larson (1981) found that the velocity dispersions of clouds as measured from the line widths σ are proportional to the size of the region observed L

$$\sigma \propto L^{\beta}$$

where $\beta \sim 0.5$. Thus the larger the region that is observed, the larger the velocity dispersion. This relationship has been ascribed to turbulence, which has become a major factor in modern theories of star formation.

Turbulence is a rather complex subject with a huge literature on both terrestrial and astrophysical turbulence. Turbulence is a process by which energy on one scale can be transferred to smaller scales through a series of "eddies" where an eddy is a local vorticity.[3] Turbulence proceeds through a "cascade" where energy on large scales can be transferred to smaller scales,

[3]Energy is transferred through vortex stretching in which local velocity gradients can amplify the vorticity.

and then to smaller scales still through eddies. Eddies are not isolated, larger eddies will contain smaller eddies, which can contain still smaller eddies.

Importantly, turbulence is a dissipative process. Energy is transferred from the largest scale downward and eventually the scale will be such that the energy is able to dissipate. At the smallest scales, turbulence will dissipate due to molecular viscosity, although in astrophysical turbulence the scale on which self-gravity dominates adds a somewhat larger scale on which the behavior of turbulence changes.

In supersonic turbulence energy is also dissipated by shocks. Supersonic collisions cause shocks which are discontinuities in the gas properties, usually seen as a strong density jump across the shock. The strong enhancement of density in shocks provides the initial conditions for self-gravitating fragments to form which will eventually become stars.

Turbulence can be described by examining the relationship between the wavenumber (scale^{-1}) κ of the turbulence, and the energy contained on that scale $E(\kappa)$

$$E(\kappa) \propto \kappa^{n}$$

where n measures the spectrum of the turbulence. For pressureless turbulence, the turbulence is expected to have the Kolmogorov–Burgers spectrum of $n = 1.75$ (Boldyrev 2002).

It is hopefully obvious that if there is more energy on larger scales, the observed velocity dispersion would increase with the size of the region observed as seen by Larson (1981). Indeed, the linewidth-size relation with $\beta \sim 0.5$ that is observed in MCs is what would be expected from a Kolmogorov–Burgers spectrum (McKee and Ostriker 2007).

2.2.6 The Formation of Cores in Molecular Clouds

In shocks and clumps the increased density can form gravitationally bound prestellar cores. Prestellar cores are the basic unit of star formation as they will each produce an individual stellar system (a single star, or often a binary or multiple system) typically with a size of <1, 000 AU. Prestellar cores may form in relative isolation (such as Bok Globules), or in their hundreds of thousands to make massive clusters.

An object will collapse if its mass exceeds the local Jeans mass. The Jeans mass M_J is given by

$$M_J = \frac{\pi}{6} \frac{c_s^3}{G^{3/2} \rho 1/2} \tag{5.1}$$

where c_s is the sound speed of the gas, ρ is its density, and G is the gravitational constant. For a typical molecular cloud with a temperature of around 10 K the sound speed is ~0.2 km s^{-1}, and the typical number density in a clump is $n = 10^3$–10^4 cm^{-3}and so the Jeans mass can be rewritten in more useful units as

$$M_J \sim (2\mathrm{M}_\odot) \left(\frac{c_s}{0.2\,\mathrm{km\,s^{-1}}} \right)^3 \left(\frac{n}{10^3\,\mathrm{cm^{-3}}} \right)^{-1/2}$$

Therefore the Jeans mass in a typical molecular clump in a cloud is roughly a solar mass. This is remarkably close to the mean mass of a star of ~0.4 $\mathrm{M}_\odot$.

A prestellar core can be created if around 1 $\mathrm{M}_\odot$of gas can be compressed to $n > 10^3$ cm^{-3}which shocks from supersonic turbulence can easily manage. It should be noted for later that to make a core of 0.1 $\mathrm{M}_\odot$the density needs to be roughly 100 times greater, and that cores of 100 $\mathrm{M}_\odot$will typically contain 100 Jeans masses of material.

2.2.7 The Formation of Molecular Clouds

The mechanism by which MCs form is unclear. Opinion is split on whether MCs are long-lived and virialized (Mouschovias et al. 2006; Tan et al. 2006) or transient, nonequilibrium structures (Ballesteros-Paredes 2006; Elmegreen 2000). This has important consequences as to the timescale of star formation: Is it long-lived and quasi-static, or rapid and dynamic? The fundamental question is whether star formation occurs on only one or two crossing times (rapid) or many crossing times (quasi-static). (See also reviews by Mac Low and Klessen 2004, and Ballesteros-Paredes et al. 2007).

Because turbulence is dissipative (see above), it does not last forever. Turbulence will decay on roughly the crossing time of the system, where the crossing time, t_{cross}, of the system is

$$t_{\text{cross}} \sim L/c_s.$$

where c_s is the (kinetic) sound speed, typically ~0.2 km s^{-1}.

Therefore if supersonic turbulence is observed it means that either (a) the MC is less than a few crossing times old, or (b) the turbulence is being "driven" (i.e., energy is being added to the MC) to maintain it. Which of these is occurring is unclear, and the timescale of star formation is somewhere between one and ten crossing times (roughly between 1 and 10 Myr for a typical pc-scale clump).

An obvious way to determine the timescale of star formation is to measure the age-spread of stars within a cluster. However, this is extremely problematic. For example, Palla et al. (2005) claim a 10 Myr age-spread amongst stars in the Orion Nebula Cluster. Given that the Orion Nebula Cluster is only about 1 pc in size with a crossing time of around 1 Myr this would seem to argue for quasi-static star formation. However, Burningham et al. (2005) and Mayne and Naylor (2008) show that at least some (but possibly not all) of the age-spread may be accounted for by binarity, variability, and photometric errors. Rather worryingly, Naylor (2009) suggests that the ages of young clusters have been under-estimated by a factor of 2 due to uncertainties in PMS tracks.

Another clue might be from the formation mechanism of MCs. If the initial conditions of MCs are known, then how star formation will then proceed in the cloud may become clearer. Unfortunately, the formation mechanism(s) of MCs are unknown. It is thought that dense molecular gas forms when the warm neutral medium that makes up the majority of the volume of a galaxy is compressed or overrun by a shock; however, many possible mechanisms for this exist and there is little agreement on which are most important (see, e.g., Vázquez-Semadeni et al. 2007 and references therein).

2.3 Low-Mass Star Formation

Once a *bound* prestellar core has formed within a MC it will collapse to form a star or stars.

Before proceeding it is useful to define the two classes of young stars: protostars and pre-main sequence (PMS) stars. Usually "protostar" is used for very young stars before they have collapsed to stellar densities when they have radii of order 1 AU. A "PMS star" is a star that has collapsed to stellar densities but has not yet started to produce energy through hydrogen fusion at its center.

2.3.1 The Physics of Core Collapse

As described above, star formation typically begins with a prestellar core of mass ~1 $M_\odot$ at a temperature of ~10 K. Initially, the H_2 in the core loses energy by being in thermal balance with the dust. Impacts with dust grains transfers thermal energy from the gas to the dust which is then able to radiate that energy away as a blackbody. At the typical densities of a prestellar core of ~10^{-20} g cm^{-3} the core is optically thin to this radiation with a peak wavelength of ~200 μm and the core temperature remains constant.

As the core collapses it is initially able to retain this thermal balance and the dust is able to radiate away the kinetic energy released by the collapse and so the temperature remains at ~10 K and the collapse is isothermal.

At a critical density of $\rho_{\rm crit} \sim 10^{-13}$ g cm^{-3} the combination of accelerating release of gravitational energy and the increasing column density of the gas mean that it is no longer able to radiate away the heat of the collapse efficiently. At this point the core begins to increase in temperature and the collapse becomes adiabatic, and it is around this point that the first protostar(s) will form. Protostars initially have a size of order 1 AU and remain at this size for some time as they are only able to radiate on a Kelvin–Helmholtz timescale. This situation continues as protostars accrete more material and increase in temperature.

The thermal behavior of gas in a core can be simply described by a barytropic equation of state of the form

$$\frac{P(\rho)}{\rho} \equiv c_s^2(\rho) = c_0^2\left[1 + \left(\frac{\rho}{\rho_{\rm crit}}\right)^{1-\gamma}\right] \qquad (5.2)$$

where P is the pressure, ρ the density, c_s is the general isothermal sound speed, and $c_0 \sim$ 0.2 km s^{-1} is the isothermal sound speed in low-density gas. γ is the adiabatic index which is 5/3 for a monotomic gas, and 7/5 for a diatomic gas. Figure 5-3 shows the temperature–density relationship for a barytropic equation of state with γ = 5/3 and 7/5: an isothermal collapse followed by adiabatic heating.

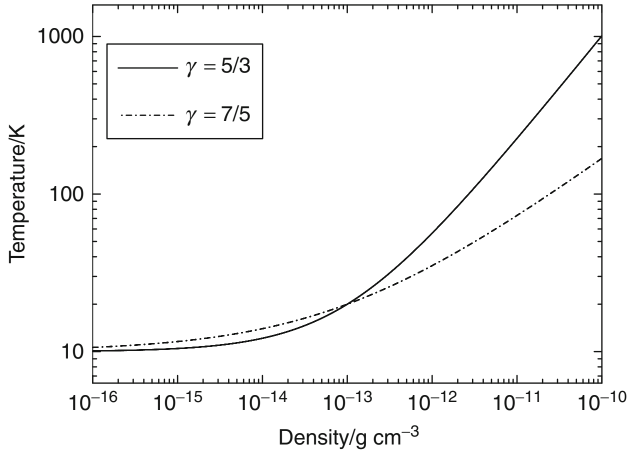


Fig. 5-3
The dependence of temperature with density during core collapse for a barytropic equation of state (5.2) with adiabatic index γ = 5/3 and 7/5

A consequence of this thermal behavior is that the Jeans mass of the core changes in an unusual way. A core is supported against collapse by thermal energy (c_s), and encouraged to collapse by its mass/density (ρ). The effect of the thermal behavior of the core is to lower the Jeans mass during isothermal collapse (ρ increases while c_s remains constant), before increasing it during adiabatic collapse (as the c_s^3 term dominates over $\rho^{-1/2}$). This means that there is a minimum Jeans mass during core collapse that is reached at a density of around 10^{-13} g cm^{-3} of around 10^{-2} $M_\odot$ (roughly ten Jupiter masses). ➋ *Figure 5-4* shows the behavior of the Jeans mass with density for barytropic equation of state with $\gamma = 5/3$ and $7/5$. The Jeans mass has a lower minimum when $\gamma = 7/5$ as the temperature rises more slowly with density. The minimum Jeans mass that is reached during core collapse is known as the *opacity limit for fragmentation* and is the minimum mass that any object can have if it forms by gravitational fragmentation.

Note that there is another minimum during the second collapse as molecular hydrogen dissociates which is potentially lower; however, it is unclear if fragmentation can occur during this phase, and it would only produce objects within 1 AU of each other. It could possibly explain some close binaries.

This situation continues until the temperature of the protostar reaches ~2,000 K at which point enough energy is available to dissociate molecular hydrogen. This dissociation provides a heat sink for the protostar and it rapidly (and almost isothermally) collapses to stellar densities (approaching 1 g cm^{-3}).

After all of the molecular hydrogen has been dissociated, the (PMS) star then slowly contracts again until the central density and temperature are high enough to start hydrogen fusion and the star joins the main sequence.

It is an interesting coincidence that the opacity limit for fragmentation of ~10^{-2} $M_\odot$ is also the point at which the planet-brown dwarf distinction is drawn. The planet-brown dwarf limit is usually taken to be the deuterium burning limit; however, the opacity limit for fragmentation might provide a far more physical distinction between planets and brown dwarfs – planets form by core accretion, while brown dwarfs form by gravitational fragmentation. The lack of

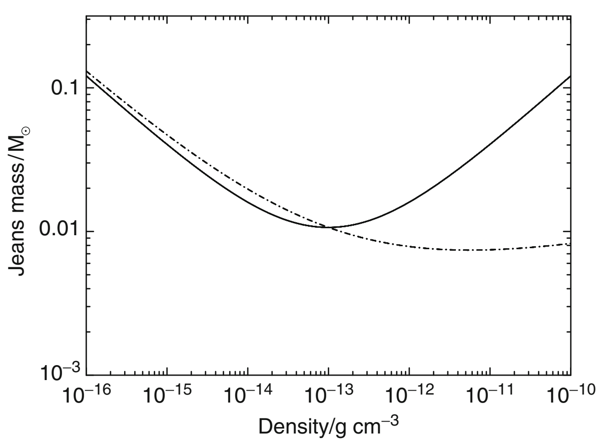


Fig. 5-4
The dependence of the Jeans mass with density during core collapse for a barytropic equation of state (➋ 5.2) with adiabatic index γ = 5/3 and 7/5

objects around ten Jupiter masses is explained by the difficulty of building a planet to such large masses, and the difficulty of creating a brown dwarf right at the lower limit of gravitational fragmentation.

For more detailed descriptions of the thermodynamics of core collapse see Larson (1969) or Masunaga and Inutsuka (2000).

2.3.2 The Stages of Star Formation

The evolution of a star forming core and then a young star is divided into four stages (classes). Figure 5-5 illustrates schematically how the spectral energy distribution (SED) of a low-mass core and/or star evolves during star formation (from Lada 1999).

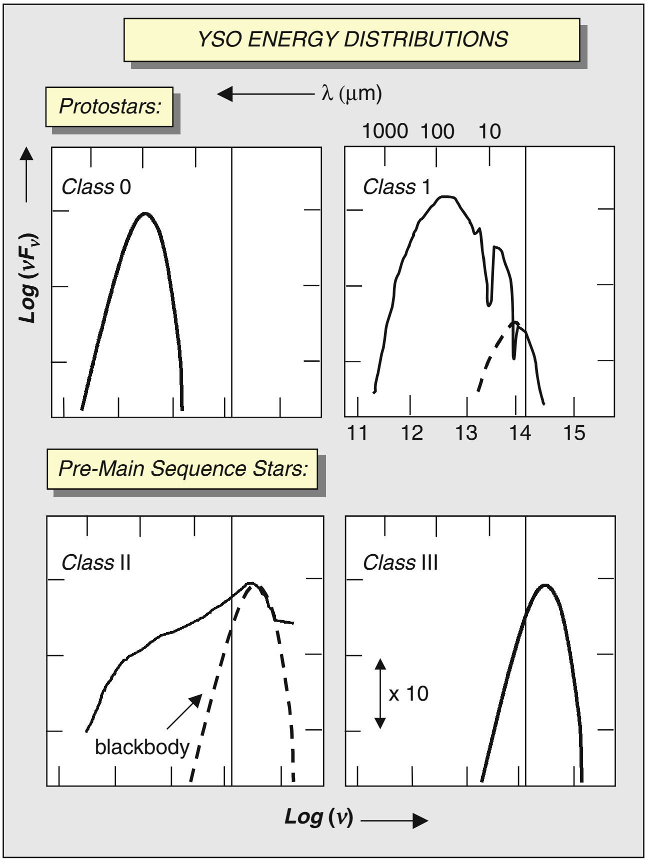


Fig. 5-5
The evolution of the spectral energy distribution of young stellar objects (YSOs) through the class 0 to class III phases of star formation (Figure from Lada 1999). Each tick on the x-axis represents an order of magnitude increase in energy emitted

Class 0. The very earliest stage of star formation is the almost free-fall collapse of the prestellar core and the formation of a class 0 object. In this stage the center of the core forms a protostar, but the majority of the mass of the core is still in the envelope. As the protostar is still heavily embedded in its natal envelope, it is optically invisible and is usually only seen as an IR point source. The emission from a class 0 core is a blackbody from to cold (few 10s K) dust with a peak in the sub-mm at around 100 μ due to the low gas temperature.

During the class 0 phase jets and outflows may also be visible. The protostar will have a strong magnetic field that is able to launch a small amount of material at high velocities along open magnetic field lines. This creates strong bipolar outflows which may be visible in a class 0 source. (The details of jet formation are extremely complex and well beyond the remit of this chapter, see the chapter on PMS stars for details).

The lifetime of the class 0 phase is extremely short as it occurs on roughly a free-fall time of the prestellar core which is $<10^5$ years.

Class I. The embedded protostar continues to accrete material from the envelope and also develop a massive disk. At the same time the protostar is heating the surrounding envelope by radiating the gravitational potential energy released by the collapse as well as potentially very significant energy input from increasingly prominent jets and outflows.

Once around half of the initial core mass has been accreted onto the protostar (and its massive disk) it enters the class I phase. Observationally the start of the class I phase is when the temperature of the envelope reaches 70 K as the protostar is still embedded and optically invisible. In the second panel of ➋ *Fig. 5-5*, the dashed line at the bottom right shows the contribution of radiation directly from the star to the SED, but most of the radiation is reprocessed by the envelope with a peak at a few tens of microns.

The class I phase lasts a few $\times10^5$ years, maybe 10^6 years, and usually it is at some point in the class I phase that the protostar collapses down to become a PMS star (this depends on the protostellar mass).

Class II/Classical T Tauri. By the end of the class I phase the bulk of the envelope has been accreted onto the (now) PMS star and its disk. The class II phase is now optically visible and is commonly known as a Classical T Tauri star (CTTs, named after the first object of its type T Tauri in the Taurus star forming region). As can be seen in the third panel of ➋ *Fig. 5-5*, the dominant contributor to the SED is radiation directly from the (PMS) stellar photosphere. However, CTT stars show evidence for their disks through an IR excess where radiation from the star is reprocessed by the cool disk and reradiated creating an excess of long wavelength radiation as seen in ➋ *Fig. 5-5*. As CTT stars continue to accrete material from the disk onto the star they show strong Hα and X-ray emission as that material collides with the surface of the star in an accretion shock. CTT stars also have strong jets and magnetically driven outflows.

The class II/CTT phase lasts for a few $\times10^6$ years until the disk is largely depleted (by a combination of accretion onto the star, planet formation, and evaporation by stellar radiation).

Class III/Weak-Lined T Tauri. Once the disk is depleted of gas the strong signatures of accretion and outflow that characterize the CTT phase also disappear. In the weak-lined T Tauri (WLTT) phase only very weak spectral signatures and a slight IR excess from the disk remain. This phase is observationally difficult to distinguish from a main sequence star except that the star is rather over-luminous for its color as it has a larger radius than a main sequence star of the same mass as it is still contracting. During this phase the rest of the disk dissipates (leaving only a debris disk in some cases). Depending on the mass of the star this phase can take a few $\times10^7$ years.

2.3.3 Different Types of Young Stars

The most common type of PMS star is a (classical or weak-lined) T Tauri star which is a typical PMS phase of stars <2 $M_\odot$, that is, the typical PMS phase of well over 90% of stars. But there are other classes of PMS stars.

Some <2 $M_\odot$ stars are classed as FU Orionis stars which are variable, with outbursts leading to increases of up to six magnitudes in a few months. Such stars are thought to be a subclass of T Tauri stars in which the accretion rate onto the star increases hugely. Typical T Tauri stars have accretion rates of 10^{-8} $M_\odot$ year^{-1}, while FU Orionis stars are estimated to be as high as 10^{-4} $M_\odot$ year^{-1}. With such high accretion rates FU Orionis stars cannot last very long (as a 0.1 $M_\odot$ disk would be depleted in only 10^3 years). It may be that many T Tauri stars undergo a very short-lived FU Orionis phase at some point in their lives, or that very few do, but that they last longer and they rapidly deplete their disks.

The PMS phase of stars of 2–8 $M_\odot$ are known as Herbig Ae/Be stars. They share most of the properties of T Tauri stars: IR excess (disks), Balmer emission, X-ray emission, and often variability.

The PMS phases of stars more massive than about 8 $M_\odot$ are not observed as such massive stars are able to reach the main sequence while still in the deeply embedded (class 0/I) phase. The formation of massive stars will be discussed in more detail later.

2.4 The Initial Mass Function

The initial mass function (IMF) is the distribution function of individual stellar masses at birth. There are many ways of parameterizing the IMF, probably the most popular (certainly theoretically) is the Kroupa (2002) three-part power-law illustrated in ➲ *Fig. 5-6* which has the form

$$N(M) \propto \begin{cases} M^{-\alpha_1} & m_0 < M/\mathrm{M}_\odot < m_1\,, \\ M^{-\alpha_2} & m_1 < M/\mathrm{M}_\odot < m_2\,, \\ M^{-\alpha_3} & m_2 < M/\mathrm{M}_\odot < m_3\,, \end{cases}$$

This parameterization usually divides the IMF into three regions.

Firstly, a low-mass (substellar) regime between 0.02 and 0.08 or 0.1 $M_\odot$ with declining numbers of stars with mass with a slope of $\alpha_1 = 0.3$.

Secondly, a roughly flat intermediate-mass region which covers M-dwarfs and the bulk of stars by a number between around 0.1 and 0.5 $M_\odot$ with a slope of $\alpha_2 = 1.3$.

Thirdly, a high-mass regime in which the numbers of stars by mass decline with a slope of $\alpha_3 = 2.3$, also known as the Salpeter slope after the first determination of the IMF by Salpeter (1955).

It should be noted that recently Chabrier (2003) has proposed a log-normal-like parameterization which is becoming increasingly popular as well.

The number of stars by mass drops rapidly (seemingly as a power-law), and extremely high-mass stars >10 $M_\odot$ which become core collapse supernovae are very rare. Below ~0.08 $M_\odot$ objects are too small to burn hydrogen in their cores (brown dwarfs). While brown dwarfs are relatively numerous (one for every five or six stars), there is again a rapid decline in their numbers with decreasing mass, and they only contribute negligibly to the total mass in stars.

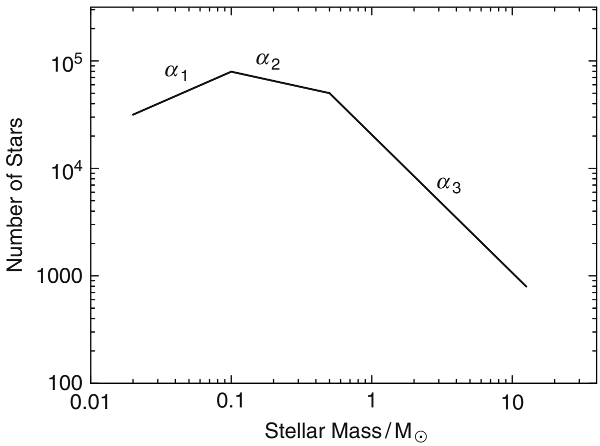


Fig. 5-6
A schematic representation of the IMF as a three-part power-law (Kroupa 2002). The three regions $-\alpha_1 = -0.3$ when $0.02 < M/M_\odot < 0.1$, $-\alpha_2 = -1.3$ when $0.1 < M/M_\odot < 0.5$, and $-\alpha_3 = -2.3$ (the Salpeter 1955 slope) when $M > 0.5\,M_\odot$

Before proceeding it is worth emphasizing two points about the IMF.

Firstly, the IMF is the *initial* mass function which is the mass function that stars have at birth. Stellar evolution (such as the supernovae of the most massive stars) will remove massive stars over time, and dynamical evolution will tend to eject low-mass stars from star clusters. Both of these processes turn the IMF into a present day mass function (PDMF) which may be very different from the original IMF. But different PDMFs do not necessarily mean different IMFs.

For example, old globular clusters and very young star clusters have very different mass functions (e.g., old globulars completely lack stars larger than ~1 $M_\odot$ due to stellar evolution). In old globulars observations are of a population that has had 12 Gyr of stellar and dynamical evolution, while in very young clusters an unevolved IMF is seen. But simulations of the evolution of globular clusters suggest that to have the PDMF that is observed today, they must have started with an IMF very similar to that of very young star clusters (Vesperini and Heggie 1997).

Secondly, the IMF is the mass function of *individual* stars. Many stars form in binary and multiple systems and the construction of a true IMF involves correcting for these multiple systems. A G-dwarf-M-dwarf binary should contribute a single G-dwarf and a single M-dwarf to the IMF. In practice this is often impossible to achieve as many multiple systems may not be known to be multiple systems (especially if the companions are of low-mass and relatively close to the primary star) and so cannot be included as two systems. Therefore what is often presented as an IMF is often a primary star IMF (i.e., many companions are missed), or a system IMF (i.e., companion masses are merged with the primary mass).

These problems notwithstanding, the IMF is an extremely useful tool for examining star formation.

The form of the IMF suggests that the vast majority (90%) of stars are M-dwarfs, with the mean mass of a star being ~0.4 $M_\odot$, and the median mass of a star being 0.2–0.3 $M_\odot$.

Surprisingly, the IMF seems to be remarkably universal. Wherever we observe the same basic form to the IMF is seen – few brown dwarfs, a peak at 0.2–0.3 $M_\odot$, and a decline to higher masses with a slope of roughly 2.3 (Salpeter). Kroupa (2002) compiled a large database of IMF (and PDMF) determinations for clusters and regions in the Milky Way and LMC of different ages and metallicities and found no significant variations from a "standard" IMF. More recently Bastian et al. (2010) have examined many claims for nonstandard IMFs and found no strong evidence for significant deviations from a universal IMF (often while a different IMF can explain unusual observations, many other more reasonable explanations can be found such as different extinction laws).

2.4.1 The Origin of the IMF

There is no a priori reason why stars forming in 10^6 $M_\odot$ proto-globular clusters with metallicities of only 10^{-3} solar should have the same IMF as stars forming in 10^2 $M_\odot$ loose associations with solar metallicity. Indeed, it might be expected that the IMFs *should* be different, the physics of star formation appear to depend on the thermodynamics of the gas and the opacity limit for fragmentation which should depend on metallicity and density. So what is the origin of the IMF and why is it apparently universal? For more details the reader is directed to Bonnell et al. (2007) and Bastian et al. (2010).

2.4.2 The IMF from Cores

Stars are seen to form in dense molecular cores. Presumably the mass of the core and the mass of the star (or stars) that form in them is related to the mass of the core; therefore, the distribution of stellar masses could well be related to the distribution of core masses. In such a scenario, the universality of the IMF may well just reflect the universality of the core mass function (CMF).

This picture is supported by observations of the CMF in different regions where the CMF is seen to have an IMF-like shape. In *Fig. 5-7* the log-normal-like CMF in the Orion B region is shown (as observed by Nutter and Ward-Thompson 2007) compared to a standard Kroupa IMF which has been shifted in mass by a factor of 8 (Alves et al. 2007 find a very similar result in the Pipe nebula). Quite clearly the shapes of the two mass functions are very similar, with the CMF just shifted in mass relative to the IMF. This shift might be reflecting that each core does not just produce one star, and that not all of the mass in the core ends-up on a star (due to feedback from jets and outflows). Therefore it might seem reasonable to conclude that each core forms two or three stars with only around 50% of the mass in the core eventually being accreted by the stars and so the form of the IMF is just set by the form of the CMF (Goodwin et al. 2008).

In many ways this may seem to be just shifting the question of the origin and universality of the IMF back one stage to the question of the origin and universality of the CMF. However, it is probably far less surprising that the CMF is universal as its origin lies in the nature of turbulence in MCs.

As described earlier, MCs are dominated by supersonic turbulence with apparently a universal power spectrum. Dense cores are formed in colliding/converging regions within MCs and the mass spectrum of cores will depend on the power spectrum of the turbulence and its strength. A core will collapse to form a star if it exceeds the local Jeans mass. The Jeans mass in a MC is ~1 $M_\odot$ at a density of about 10^{-18} g cm^{-3}, and so to form a prestellar core of 1 $M_\odot$ turbulence needs to compress 1 $M_\odot$ of gas to a density of 10^{-18} g cm^{-3} – only a compression of a factor of about 10^3 or 10^4 above the average density in a MC. However, to make a

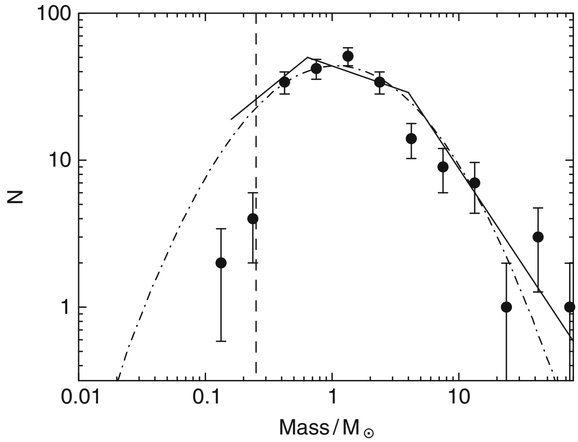


Fig. 5-7
Points show the core mass function of Orion B (from Nutter and Ward-Thompson 2007) compared to a Gaussian (*dot-dashed line*), and a Kroupa (2002) IMF shifted in mass by a factor of 8 (*solid line*). The *vertical dashed line* shows the completeness limits of the observations (Figure from Goodwin et al. 2008)

0.1 $M_\odot$ prestellar core the compression required is an extra factor of 100. Much larger cores, even though they require lower densities will be far larger in extent and expected to fragment into smaller cores unless they start at relatively higher densities (i.e., very super-Jeans).

Therefore very low-mass cores are expected to be rare as they require very high compression to reach the densities required for them to collapse, while cores of ~1 $M_\odot$ should be typical, and there should be few high-mass cores as they will fragment into smaller cores unless significantly compressed. This is exactly the form seen in observed CMFs and in the IMF (see Hennebelle and Chabrier 2008; Padoan and Nordlund 2002 for a far more detailed discussion of the creation of CMFs from turbulence). However, as is discussed later, the picture is far more complex than this.

2.4.3 Competitive Accretion

Competitive accretion (Bonnell et al. 1997, 2001, 2007) is the main competitor to core formation as the origin of the IMF. In competitive accretion stars form in cores which follow some CMF, but it is not the initial mass of the core that is important in setting the final mass of the star. Rather, cores, and later stars, continue to accrete gas as long as there is gas in the cluster to accrete (i.e., before it is blown away by feedback from massive stars).

Thus there is no such thing as a "core mass" which sets the mass of stars which form within it. The observed CMF is just the instantaneous amount of gas in dense condensations. Many of these condensations will collapse and form a star following the classical class 0 through class III evolutionary path. However, some stars will continue to accrete gas and continue to grow.

A key element of competitive accretion is that not all stars are equally successful in accreting material. More massive stars will be more successful in accreting material due to their larger gravitational influence, but more importantly some stars will just be "lucky" in that they happen to be in regions where there is a lot of gas to accrete, while other stars will be in "unlucky" regions. Thus what becomes a massive star has no need to start as a particularly massive star. The protostar which forms the most massive star eventually can begin life in a small core.

2.4.4 The CMF Verus Competitive Accretion

Given that CMFs are observed to have a form that is very similar to the IMF it might be thought that this scenario is the most natural. However, there are a number of problems associated with the CMF-to-IMF model.

Firstly, the peak and width of the CMF should depend strongly on the Mach number of the turbulence. More energy in turbulence means stronger compression and a lower peak and more low-mass cores. Therefore the IMF should depend on the Mach number (Hennebelle and Chabrier 2008; Padoan and Nordlund 2002). MCs with Mach numbers ranging between around 5 and 20 have been observed (Ballesteros-Paredes et al. 2007), but the IMF is universal.

Secondly, it is uncertain how the observed CMF is related to the final IMF. It is often not clear which cores will collapse to form a star and which are transient density enhancements (i.e., which cores are really prestellar and which are not). In addition, cores of different masses evolve at different rates, and so any snapshot observation of a CMF is not a snapshot of the full range of cores that will eventually form stars (Clark et al. 2007). In addition, cores are often difficult to identify, and different methods can find different cores and different CMFs in the same data (Smith et al. 2008).

Indeed, it is not clear that the observed cores are the precursors of many stars. The typical cores that are observed have a size of roughly 0.1 pc. However, most stars form in dense clusters which contain (to take the example of the well-studied Orion Nebula Cluster) about 10^3 stars in a cubic pc. At such densities, the average distance between stars is the size of a core and so the filling factor of cores in the proto–Orion Nebula Cluster must have been unity which seems unlikely (Goodwin et al. 2007).

Thus the cores which form stars in dense clusters must have significantly smaller sizes, but the same mass spectrum as the more isolated cores with which CMFs are constructed. There is clearly an observational bias here, as the sizes of cores that can be observed are set by the resolution limits of the instruments used to observe them. There are no dense proto-clusters close enough to resolve many very small cores, and in more distant dense proto-clusters small cores would be unresolved. Indeed, it is found that the typical core mass increases with distance (resolution) which suggests that observers are unable to resolve smaller structures in these larger "cores." It may be that all of the substructure cannot be resolved, even in nearby prestellar cores, and that they may contain substructure below current resolution.

It is likely that both the CMF *and* competitive accretion play a role in star formation and establishing the universal form of the IMF. Given the apparent link between the CMF and the IMF in diffuse regions where competitive accretion is likely to be less effective, it would appear that the CMF does largely set the form of the IMF. However, in dense environments it is difficult to imagine that cores and protostars could avoid accreting at least some of the large amounts of ambient gas around them.

2.5 The Formation of Massive Stars

The typical mass of a star of about 0.4 $M_\odot$ is probably set by the typical mass of a core of about 1 $M_\odot$, which is itself probably determined by the Jeans mass in a MC. However, the origin of stars significantly more massive than 1 $M_\odot$ is very uncertain (see Zinnecker and Yorke 2007 for a detailed review of massive star formation).

There is a major theoretical problem in forming stars in excess of ~10 $M_\odot$, despite observations of many stars that appear to have masses well in excess of 100 $M_\odot$, possibly up to 300 $M_\odot$. As soon as a star reaches a mass of ~10 $M_\odot$ its central temperature and pressure will be high enough to begin core hydrogen burning. More massive stars evolve at a much faster rate than low-mass stars, a star >40 $M_\odot$ will evolve through the main sequence and post-main sequence to become a core collapse supernova in only a few Myr. And during their rapid evolution they have extremely strong stellar winds and UV fluxes. The question is then, if a star of ~10 $M_\odot$ is producing strong winds and a strong ionizing UV field, how does more material accrete onto the star to produce a star of 20, 50, or even 300 $M_\odot$?

It is extremely difficult to observe young massive stars as they are always deeply embedded, and once they reach the main sequence, they produce an HII region. Their evolution can in some ways be tracked through the properties of the HII regions. Initially, a massive star will produce a hypercompact or ultracompact HII region which is less than 0.1 pc in size with densities of 10^4 cm^{-3}. As the HII region evolves it will increase in size and decrease in density toward a classical HII region of pc-scale as the radius of its Strömgren sphere increases. Once a massive star (or often stars) have ionized a large region, they begin to stop star formation locally. But it is in the hyper- and ultracompact phases that massive star formation is occurring and it is extremely difficult to observe the processes that are occurring.

In order to build a massive star accretion must overcome the force of radiative (UV) and mechanical (winds) feedback. Massive stars can have luminosities of >10^5 $L_\odot$ and mass-loss rates of >10^{-5} $M_\odot$ year^{-1} which is a significant load to be overcome. By far the best way to do this seems to be via a disk. Even if feedback from the star is initially isotropic the presence of a disk means that it rapidly becomes anisotropic and escapes preferentially toward the poles allowing accretion through the disk (see Zinnecker and Yorke 2007 for details).

If it is possible to overcome feedback (and it must be as very massive stars exist), the formation of a massive star requires a large amount of gas to be available to accrete. There are two theories as to the origin of this huge reservoir of gas.

Possibly most obviously given our discussion of the core mass function is that massive stars form from massive cores (McKee and Tan 2003). Given the large amounts of feedback from a massive star, the core mass presumably needs to be significantly greater than the mass of the final star, and a 50 $M_\odot$ star will need a ≫50 $M_\odot$ core from which to form. Such massive cores are observed (as "clumps," see above), but it is very unclear why a massive core should produce only one or two stars. A 100 $M_\odot$ core contains many Jeans masses of material, and might be expected to form a small cluster rather than a single massive star. In addition, massive stars tend to be found in groups near the centers of clusters. For example, the Trapezium in the Orion Nebula Cluster contains four of the six most massive stars in the cluster in a region only around 0.1 pc across. But massive cores are by definition large, and so are not expected to form in the centers of clusters.

The alternative scenario for massive star formation is competitive accretion (see above). Here there is no need to form a massive core, rather massive stars begin as "normal" mass stars. A lucky few "normal" stars will sit in the deep potential where gas can be directed toward them from which they can accrete. Thus they need no large initial reservoir of gas as the gas is channeled toward them. This has the advantage of automatically producing massive stars toward the centers of clusters where the potential is the deepest. However, it is not clear if a constant infall of gas can be maintained through the feedback from the most massive stars. Gas must accrete onto the massive star from a disk which can channel the feedback around it. If the inflow of gas into the center of the cluster is low-density then it could well be disrupted by feedback.

2.6 The Formation of Brown Dwarfs

At the other end of the mass spectrum from massive stars are very low-mass stars and brown dwarfs, and, as with massive stars, their formation is rather difficult to explain. The most obvious way in which brown dwarfs (and very low-mass stars, typically anything with a mass $<0.1\,M_\odot$) is in low-mass cores in a very similar way to stars. Young brown dwarfs appear very similar to young stars (Whitworth et al. 2007), and so their formation mechanism might well be exactly the same.

However, it may be difficult to form cores of a low enough mass to form brown dwarfs. To become bound a $0.1\,M_\odot$core must be 100 times denser than a $1\,M_\odot$core (due to the $\rho^{-1/2}$ dependency of the Jeans mass). This implies that very low-mass cores should only form where there is a strong shock which is able to increase the local density significantly. In particular, it implies that low-mass cores should form when the turbulence is strong (high Mach-number turbulence), and not when it is relatively weak. A problem arises in that the substellar IMF appears to be the same in all regions, even in low-density, relatively low-turbulence star forming environments such as Taurus.

In addition, very few very low-mass cores are observed. But for every very low-mass core which is able to become bound around 10^4 should be formed which just fail to reach the required densities and then disperse (Goodwin and Whitworth 2007). Therefore, huge numbers of very low-mass cores should be observed and they are not.

An alternative is to form brown dwarfs in higher-mass cores which we know to exist in significant numbers. However, the problem is that once an object has formed at around the opacity limit for fragmentation of $\sim 10^{-2}\,M_\odot$if there is a significant amount of gas present, then it will accrete that gas (as happens in a "normal" star). Therefore some mechanism is required to stop the object accreting and to keep it at brown dwarf masses.

The first suggestion to accomplish this was the ejection scenario (Reipurth and Clarke 2001). In this scenario several very low-mass objects form at roughly the same time in a core. Systems with $N > 2$ are generally unstable and will rapidly decay and will usually eject the lowest-mass member of the system. Thus, brown dwarfs can be ejected from cores shortly after their formation and once they have left the core they have no more gas available for them to accrete.

The problem with this model is that brown dwarfs are typically ejected with speeds of a few km s^{-1}. This means that in a Myr an ejected brown dwarf could travel a few pc, and so clusters should have a halo of ejected brown dwarfs and the distribution of brown dwarfs should be different to that of stars. However, brown dwarfs and stars appear to have the same spacial distribution which suggests that violent ejection cannot be the answer (Luhman 2004).

Other possibilities have been suggested, such as forming brown dwarfs in massive extended disks around solar-type stars which can then be gently liberated (Stamatellos et al. 2007), or that brown dwarfs form as wide companions to M-dwarfs which are then disrupted

(Goodwin and Whitworth 2007). But every formation model has some problems, and the formation mechanism of brown dwarfs remains uncertain (indeed, they may form from a combination of all the processes that have been suggested).

2.7 Star Formation Efficiency

How much of its gas does a MC convert into stars? This may appear to be a simple question to answer: simply observe the total mass of gas M_g in a MC and the total mass of stars $M_\star$, and the star formation efficiency (SFE) is then SFE $= M_\star/M_g$. However, in practice this is rather difficult to do.

In order to measure the total mass of gas in a MC, there must obviously be gas present. This means that some (many? most?) of the stars will be embedded, either in their natal core or in the general cloud making them difficult to observe, therefore it is often not clear if the observed stellar mass is the total stellar mass.

That gas is still present means that star formation is ongoing. In particular more stars will form increasing the stellar mass, and the masses of stars already present will increase. If competitive accretion is occurring in the cluster the mass in stars could increase very significantly. Therefore, even if the current mass in stars is known exactly, the SFE is an instantaneous value, and not the final SFE (e.g., Evans et al. 2009). If more stars form from the same mass of gas the SFE will increase, but if more gas is channeled into the region, by turbulent flows or along filaments, for example, then the SFE could decrease.

Over whole galaxies the SFE (usually quoted as a star formation rate) can be calculated. The gas content (HI + H_2) of entire external galaxies can be observed (usually through 21 cm + CO). In addition, the *massive* star content of external galaxies can also be seen (either from HII regions/H-α emission or naked O-stars) which, with an assumption of a standard IMF, gives the mass of young stars in that galaxy. Interestingly, the SFE in H_2 appears roughly constant in spiral galaxies at a few % (Leroy et al. 2008).

The SFE is observed to decrease with increasing scale. In prestellar cores the SFE is probably around 50% (Goodwin et al. 2008), in clumps 30–50% (Lada and Lada 2003), and in typical spiral galaxies as a whole it is usually a few percent (Leroy et al. 2008). Thus the majority of gas in a galaxy is not forming stars, and the majority of gas only ends up on a star when it reaches prestellar core scales of <0.1 pc.

3 Multiple Stars, Star Clusters, and the End of Star Formation

Stars do not form on their own. Many (possibly most) stars form as binary and multiple systems in which each core produces two or more stars on scales of a few hundred AU. In addition, most stellar systems (singles or multiples) form in star clusters of hundreds to millions of members on scales of roughly a pc.

3.1 Binary and Multiple Systems

In the discussions previously it has tended to be implicitly assumed that a single core produces a single star. However, it is known that this simplification is certainly not always true, and maybe wrong most of the time.

Observations of old main sequence stars in the solar neighborhood show that many stars are in binary and multiple systems. The numbers of stars in binaries is usually quantified by the binary fraction f_{bin}

$$f_{\text{bin}} = \frac{\text{B} + \text{T} + \text{Q} + \ldots}{\text{S} + \text{B} + \text{T} + \text{Q} + \ldots} \tag{5.3}$$

where S, B, T, Q, etc., are the numbers of single, binary, triple, quadruple, etc., *systems* (a single system has one star, a binary system contains two stars, etc.,). (There are numerous different ways of quantifying the binary fraction, see Reipurth and Zinnecker 1993 for a detailed description of many.)

For massive >5 $M_\odot$ stars the binary fraction in the field appears to be unity (i.e., all massive stars are in a binary or multiple system). For solar-type stars the fraction appears to be ~60% (Duquennoy and Mayor 1991), falling to around 40% for M-dwarfs (Fischer and Marcy 1992; see also Lada 2006), and 15% for brown dwarfs (Burgasser et al. 2007). It should be noted though that these numbers are probably lower limits as it is generally very difficult to detect companions with a significantly lower mass than the primary star.

It seems to be impossible to make binaries in the numbers seen by dynamically combining initially single stars (Kroupa and Burkert 2001) and so the *vast* majority of the binary stars must have formed as binaries. In fact, the situation is even more extreme for young stars. *Figure 5-8* shows the fraction of solar-type stars with companions at different separations for the field (solid line), and for a selection of young stars in different regions compiled by Patience et al. (2002). There are clearly far more binaries with separations of around 100 AU amongst young stars than in the field. Indeed, the binary fraction of young solar-type stars appears to be close to 100%.

This suggests that most stars, at least >1 $M_\odot$ (the situation is unclear for lower-mass stars), form as binaries.

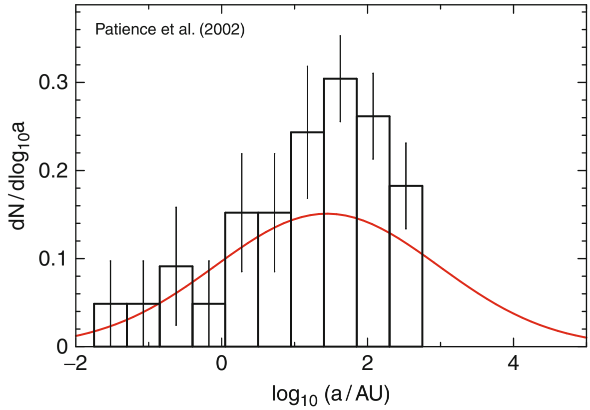


Fig. 5-8
The binary fraction with separation of field G-dwarfs (*red line* from Duquennoy and Mayor 1991), and for young (T Tauri) stars (Compiled by Patience et al. 2002)

3.1.1 Binary Formation

If binary and multiple formation is an extremely common (possibly the major) mode of star formation then it is important to ask how they form. It is thought that multiple systems form by the fragmentation of circumstellar disks during the class 0 (possibly class I) phase of star formation. Prestellar cores are observed to have at least some angular momentum (Goodman et al. 1993), probably due to small (subsonic) levels of turbulence (Burkert and Bodenheimer 2000). Low angular momentum material will form a central protostar, but high angular momentum material will form a circumstellar disk around the protostar. If this disk is massive and cool enough it will be able to fragment, forming a multiple system (see Goodwin et al. 2007 for details).

The Toomre (1964) criteria describes if a disk is unstable to gravitational fragmentation. The Toomre Q-parameter at a radius R in a disk is given by

$$Q(R) = \frac{c(R)\,\kappa(R)}{\pi G\,\Sigma(R)}$$

where c is the isothermal sound speed, κ is the epicyclic frequency, and Σ is the surface density, all at radius R. If $Q < 1$, a disk is unstable to gravitational fragmentation (the numerator is a measure of thermal and rotational support against fragmentation, the denominator is a measure of the gravitational attraction of a region).

However, in order to form a new protostar, an unstable fragment must be able to cool on a dynamical timescale (otherwise the fragment will heat as it collapses and bounce); this is known as the Gammie (2001) criterion. Therefore the cooling time $t_{\rm cool}$ must be less than

$$t_{\rm cool} < \gamma t_{\rm orb}$$

where $t_{\rm orb}$ is the orbital timescale at whatever radius R, and $0.5 < \gamma < 2$ is the adiabatic index (see Stamatellos and Whitworth 2008 for details).

Whitworth and Stamatellos (2006) use reasonable values for disk parameters and find that these conditions are only met if the radius is greater than a minimum radius $R_{\rm min}$

$$R_{\rm min} > 150 \left(\frac{M_\star}{\rm M_\odot}\right)^{1/3} \quad \rm AU$$

where $M_\star$ is the mass of the star around which the disk is attempting to fragment. The dependence of the fragmentation radius on the central stellar mass is due to the irradiation of the disk by the central star. The more massive the central star, the higher its luminosity, and so the greater the temperature of the disk.

Interestingly, this value of ~150 AU is very close to the peak of the T Tauri binary separation distribution. However, it does suggest that binaries should not *form* within $R_{\rm min}$. However, inspection of *Fig. 5-8* shows immediately that there are many binary systems with companions at distances $<R_{\rm min}$. The critical point is that companions should not form below $R_{\rm min}$, but they may later move to smaller radii due to energy conservation during an ejection if there are more than two bodies in the system (although this should not happen too much, Goodwin and Kroupa 2005), or due to interactions with the disk analogous to planetary migration (e.g., Lin and Papaloizou 1979).

3.1.2 Binary Destruction

Observations and theory both suggest that binaries are the major mode of star formation (at least $>0.5\ M_{\odot}$, Goodwin and Kroupa 2005; Lada 2006). This raises the obvious question of why there are fewer old binaries than young binaries. At some point many young binaries with separations of >100 AU must be destroyed in order to produce the old population that observed in the solar neighborhood.

Binaries can be destroyed by dynamical interactions with single stars or other binaries. How easy it is to destroy a binary depends on two factors: the strength of an encounter, and the frequency of encounters which may destroy the binary (see Heggie 1975 and Hills 1975 for a detailed analysis).

The potential of an encounter to destroy a binary depends on the binding energy E of a binary which depends on the masses m_1 and m_2 of the components, and their separation a

$$E = -\frac{Gm_1m_2}{2a}$$

and the energy of the perturbing star which will depend on its typical mass $< m >$, and its typical velocity σ (which will be the velocity dispersion of the system). If $|E|/ < m > \sigma^2 \gg 1$ the binary is said to be "hard" and is difficult to disrupt, if $|E|/ < m > \sigma^2 \ll 1$ the binary is "soft" and easy to disrupt.

For a typical binary in a fairly typical cluster $< m > = m_1 = m_2 = 0.5\ M_{\odot}$, and $\sigma = 2$ km s^{-1}. This would place the hard-soft boundary at a separation of $a \sim 50$ AU. Therefore encounters would be expected to destroy binaries with separations $\gg$50 AU, and have no effect on binaries with separations $\ll$50 AU. In the intermediate regime around 50 AU encounters may or may not destroy binaries (this regime can only be probed by numerical simulations).

However, the encounter rate also plays a crucial role in determining binary survival. The velocity dispersion in the Galactic disk is ~30 km s^{-1}which gives a hard-soft boundary of only a few AU. But there are many binaries (in fact most) in the field with separations $\gg$10–100 AU. While these binaries are soft they are able to survive because the encounter rate in the field is so low.

The rate R at which a star will have an encounter within a distance b depends on the number density n and velocity dispersion σ of the environment

$$R \sim 30 n \sigma b^2 (1 + \Omega)$$

where Ω is the Safranov number and is a measure of gravitational focusing which increases the encounter rate. For the situations that are of interest here Ω is generally <1 and may be ignored.

For an encounter with an impact parameter of $b = 500$ AU in the Galactic field with a typical number density of 1 star pc^{-3} and velocity dispersion of ~30 km s^{-1}(~30 pc Myr^{-1}), the encounter rate is roughly once every 5.5 Gyr. However, in a cluster with 10^3 star pc^{-3} and $\sigma \sim$ 2 km s^{-1}, the encounter rate is once every 5.5 Myr. Therefore encounters that would be expected to destroy binaries with separations greater than about 100 AU are very rare in the field, but are expected to have occurred in even very young star clusters.

3.2 Binaries or Singles?

That so many stars form in binaries shows that binary formation is a major mode of stars formation. But is it the main mode of star formation? Most stars (around 90%) are M-dwarfs, and the binary fraction of M-dwarfs is low. A binary fraction of only 40% suggests that most M-dwarfs are single. However, the binary fraction is for *systems*. Of every 100 M-dwarf systems 40 are binary systems, and 60 are single. Of the 40 binary systems, each contains 2 M-dwarfs, meaning that by number 80/140 M-dwarfs are in a binary system, and 60/140 are singles. This now suggests that most M-dwarfs are in binary systems. However, the 40 binary systems presumably formed as binary systems and so only 40% of low-mass stars are known to form as binaries, while 60% are *currently* singles. If all the single field M-dwarfs formed as singles this tells us that single star formation is the major mode of star formation, but if many formed as binaries which were later destroyed then binary formation is the major mode of star formation. Sadly, observations of low-mass binary fractions in clusters cannot currently distinguish these possibilities.

3.3 Star Clusters

Around 75–90% of stars form in clusters of hundreds to millions of stars (Lada and Lada 2003). A particularly interesting observation is that the mass function of these star clusters appears to be $N(M) \propto M^{-2}$ (Lada 2010; Lada and Lada 2003). This is interesting because it means an equal mass of stars forms in each equal logarithmic mass interval. Therefore there are as many star forms in Taurus-like associations with mass ~$10^2\ M_\odot$, as in Orion-like clusters with mass ~$10^3\ M_\odot$, as in massive starburst clusters like Westerlund 1 with a mass of ~$10^5\ M_\odot$. Therefore all masses of clusters are equally important, and there is no such thing as an "average" or "typical" star cluster.

Our mental image of star clusters is set by pictures of old globular clusters, which are remarkably spherical, smooth, and dynamically relaxed. This is not, however, how star clusters form. MCs are complex, hierarchical structures, and they form stars in the same complex structures (see *Fig. 5-2*). Star clusters also appear to form highly out-of-equilibrium, with velocities well below what would be expected for virial equilibrium (see Allison et al. 2009 and references therein).

Any clumpy, out-of-equilibrium system will rapidly attempt to reach equilibrium and it appears that star clusters rapidly collapse and smooth-out their initial clumpiness. This has two important consequences when attempting to interpret observations of clusters.

Firstly, clusters change their appearance rapidly and can appear very different on timescales of just 10^5 years. While this is extremely rapid on astrophysical timescales, it is somewhat longer than the average baseline of our astronomical observations. Therefore only a single snapshot in the evolution of a star cluster is observed. Two star clusters of different ages that appear very different might well just be very similar objects at different points in their evolution. Equally, two objects that appear very similar might be seen at different points in their evolution and will result in very different objects. In addition, the evolution of star clusters occurs far faster than our ability to accurately date star clusters (at best to about a Myr), so it is impossible to tell if two young clusters are really the same or of significantly different ages.

Secondly, our observations beyond ~100 pc are only two-dimensional projections of complex, three-dimensional objects. A clumpy, hierarchical cluster can look very different from different viewing angles. This should improve with the advent of GAIA which will provide accurate distances to many objects (the 10^9 brightest objects in the sky) and also proper motions to give us detailed kinematic information. However, GAIA will not be able to observe the earliest embedded stages of star cluster formation and evolution during which much of the rapid evolution occurs.

Clusters that are initially dynamically cool and clumpy will attempt to reach an equilibrium which is virialized and smooth. Following Allison et al. (2009), if the initial virial ratio of a cluster is $Q_i = -T_i/\Omega_i$, where T_i is the initial kinetic energy and Ω_i is the initial potential energy (and $Q_i = 0.5$ is virial equilibrium), then the total energy of a cluster is

$$E = T_i + \Omega_i = (1 - Q_i)\Omega_i$$

the initial potential energy is

$$\Omega_i = -\eta_i \frac{GM^2}{R_i}$$

where M is the mass of the cluster, R_i is a characteristic radius, and η_i is a structure parameter. The cluster will virialize and erase its substructure resulting in a cluster where

$$E = \frac{\Omega_f}{2} = -\eta_f \frac{GM^2}{2R_f}$$

with a new distribution and so a new structure parameter η_f, and a new radius R_f. The degree to which the cluster has collapsed is

$$\frac{R_i}{R_f} = \frac{\eta_i}{\eta_f} 2(1 - Q_i)$$

Typical values for these parameters are an initial virial ratio of $Q_i \sim 0.3$, an initial clumpy structure parameter of $\eta_i = 1.5$, and a final smooth structure parameter $\eta_f = 0.75$ (clumpy distributions have a higher η as clumps contain more potential energy than if they were smoothed-out). This means that clusters will typically collapse by a factor of about 2.5. After collapse clusters will tend to "bounce," rapidly increasing their radii (Allison et al. 2009).

An interesting consequence of this collapse and bounce is that the dynamical timescales of clusters will change significantly during their early life. The crossing time/dynamical time[4] of a cluster can be calculated if the size R and the velocity dispersion σ of a cluster are known

$$t_{\text{cross}} = \frac{R}{\sigma}$$

The crossing time gives the shortest timescale on which the cluster can be expected to change, in particular the timescale on which dynamical interactions and the destruction of binaries will occur. From the crossing time the two-body relaxation time of the cluster can also be calculated, which is the timescale on which the velocities of stars will significantly change.

The size of a cluster is relatively simple to observe, but the velocity dispersion of a cluster is often unknown. To observe the velocity dispersion of a cluster either detailed spectroscopy of many stars (which may be contaminated by binary motions) or long-baseline proper motion determinations are required. Very often these are not available and so the velocity dispersion is estimated by assuming that the cluster is virialized. This is probably a very poor approximation

[4]For our purposes the terms are interchangeable.

in young clusters as it is found that when the actual velocity dispersion is known, many clusters appear to be out of virial equilibrium (Goodwin and Bastian 2006).

However, determinations of the crossing times of clusters are *instantaneous* values. If a cluster has undergone a collapse and a re-expansion, then R may have changed by a factor of several, and σ may well have changed as well (although probably less than R) and therefore the cluster may be many more crossing times older than the instantaneous value might suggest, and so the cluster may be far more dynamically evolved than one might think.

As an example, the current half-mass radius of Orion is 0.8 pc, and its mass is about 10^3 $M_\odot$. If Orion were virialized this would suggest a velocity dispersion of $\sigma \sim 2$ km s^{-1}, and a crossing time of about a Myr giving a dynamical age of Orion of 2 or 3 crossing times. However, Orion is not virialized, and its velocity dispersion is actually observed to be $\sigma \sim 4$ km s^{-1}, implying a crossing time of only half the virialized value. But, if the velocity dispersion of Orion is significantly super-virial, then Orion must be expanding, and so its size must have been significantly smaller in the past as well. Therefore Orion is probably tens of crossing times old, rather than the 2–3 implied by assuming it is virialized, or the 4–6 from using the actual velocity dispersion, but only the current (larger) size.

3.4 The End of Star Formation

Star formation ends when there is no (dense molecular) gas remaining from which to form stars. It is thought that the feedback of energy from massive stars – winds, radiation, and eventually supernovae – stop star formation by heating and then removing gas from the star forming region. Massive stars begin to feedback energy while they are still forming, and so the formation of the first massive star(s) marks the beginning of the end of star formation.

Stars >30–40 $M_\odot$ will feedback around 10^{47} J of energy in winds and UV radiation during their lifetimes, and another 10^{46} J when they become supernovae. Their prodigious energy output ionizes the gas around the stars preventing it from forming new stars and expels the gas from around them. But the expulsion of gas from a central cluster of massive stars can also trigger star formation in surrounding gas that is dense enough, and/or far enough to escape being ionized. Observations of young star clusters find that there are no star clusters older than 3–5 Myr which have gas associated with them (e.g., Lada and Lada 2003) suggesting that gas loss occurs at around this age.

It is also found that there are far fewer star clusters older than 10 Myr than would be expected from the number of young clusters we see (Lada and Lada 2003). ➲ *Figure 5-9* shows the observed numbers of clusters within 2 kpc of the Sun of different ages (solid line) against the expected numbers of clusters for a constant cluster formation rate (dotted line), and the expected number corrected for luminosity evolution (dashed line). If the cluster formation rate were constant, each equal logarithmic age bin should contain increasing numbers of clusters (as it covers a larger span of linear time).

The observations of clusters have some problems. Firstly, aging the youngest clusters is rather problematic, and so the differences between the first two bins should not be taken too seriously. Secondly, less massive clusters are expected to dissolve due to internal two-body effects and so cluster numbers will fall with time. However, the lack of older clusters compared to younger clusters is dramatic and inexplicable by either internal dynamics or errors in the observations. This tells us one of two things: (a) the cluster formation rate has not been constant over the past few hundred Myr, or (b) many (most) clusters are destroyed.

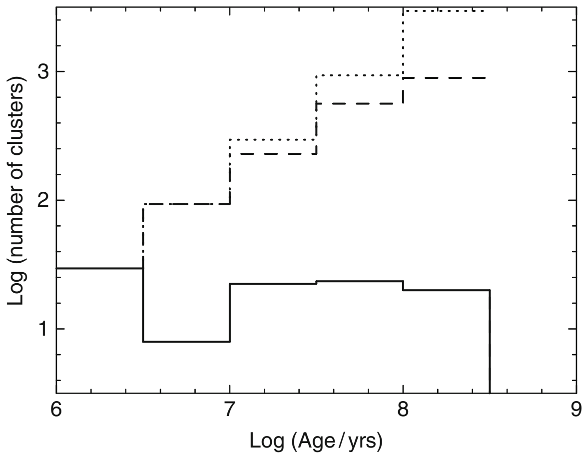


Fig. 5-9
The *solid line* shows the observed number of clusters with age within 2 kpc of the Sun. The *dotted line* shows the expected number of clusters if the cluster formation rate has been constant over this period and no clusters have been destroyed, the *dashed line* shows this number corrected for luminosity evolution (Adapted from Lada and Lada 2003)

The favored interpretation is that clusters are destroyed. There is no evidence that the current cluster formation rate is vastly (factors of 10) higher than in the recent past. In particular, there is absolutely no evidence that the cluster formation rate has vastly increased in the past 10–20 Myr as would be needed to explain the difference between the model and the second and third age bins.

Clusters are thought to be destroyed by two mechanisms. Most (up to 90%) clusters that form are thought to be destroyed by the rapid expulsion of gas by massive stars. And many of those clusters that do survive are destroyed by evaporating.

3.4.1 Gas Expulsion

As noted before, the star formation efficiencies of star clusters are low, only 10–30%. Therefore star clusters at the point at which massive stars begin to remove gas are gravitationally dominated by that gas. Therefore, the removal of the gas removes the largest contributor to the binding energy of the cluster (see Goodwin 2009 and references therein).

Assuming that a cluster of total mass M has formed stars with an efficiency ϵ, the stellar mass is ϵM and the gas mass is $(1-\epsilon)M$. For a cluster of size R and velocity dispersion σ the initial energy E_i is

$$E_i = \frac{1}{2}M\sigma^2 - \frac{GM^2}{R}$$

and if it is in virial equilibrium then

$$E_i = -\frac{GM^2}{2R}.$$

If all of the gas is expelled instantaneously (in this case, this means in less than a crossing time), then R and σ for the stars do not change, but the gas mass is removed meaning that the final energy E_f is

$$E_f = \frac{1}{2}\epsilon M\sigma^2 - \epsilon^2\frac{GM^2}{R}$$

It is easy to show that if $\epsilon < 0.5$, then the cluster is now unbound and has been destroyed by the gas expulsion.

In practice, gas expulsion is rather more complex than this as gas expulsion can be adiabatic rather than instantaneous (i.e., it takes longer than a crossing time) which somewhat alleviates its effects, and also even if it is instantaneous a much smaller bound core may be retained. But even taking these effects into account, star clusters cannot survive gas expulsion if their star formation efficiencies are less than 20–30% (Goodwin and Bastian 2006). Such star formation efficiencies are at the upper end of those observed, and so it would be expected that many star clusters would be destroyed by gas expulsion after only a few Myr in keeping with the observations of cluster numbers with age.

3.4.2 Dynamical Evolution

Even if a star cluster is able to survive gas expulsion, many will not survive for long. Two-body interactions between stars alter the velocity dispersion of the stars in a cluster driving it toward a Maxwellian. Stars in the tail of the Maxwellian with greater than the escape velocity of the cluster are able to escape, which lowers the mass of the cluster, which lowers the escape velocity, which makes escape easier. Thus a cluster will evaporate over time.

The timescale over which the velocity distribution changes significantly (and a Maxwellian is established) is the relaxation time t_{relax} given by

$$t_{\mathrm{relax}} \sim \frac{N}{8\ln N}t_{\mathrm{cross}}$$

where N is the number of stars in a cluster (see Binney and Tremaine 2008 for a derivation), and clusters can survive for roughly ten relaxation times before dissolving, although this depends strongly on the strength of the external tidal field (see Spitzer 1987). Lamers et al. (2005) provide a semi-analytic formulation of the survival time of a star cluster in a galaxy.

Therefore, low-mass (low-N) clusters will not survive for long, even if they do manage to survive gas expulsion. It is no surprise that the only extremely old clusters that are observed are very massive globular clusters, as they are the only clusters that *could* have survived for a Hubble Time, even if many low-mass clusters also formed 12 Gyr ago.

3.5 Is Star Formation Universal?

The final question is that of the universality of star formation. Do all stars form in roughly the same way, or does the environment in which stars form play a crucial role? Are the stars that form in Taurus-like associations the same as those that form in very massive clusters? And do all clusters of the same mass form stars in the same way? While this might seem like a fairly simple question, the answer is far from clear.

The form of the IMF might well be expected to vary with environment and thus provide important clues as to the universality or otherwise of star formation. However, the IMF is surprisingly invariant (Bastian et al. 2010). It is unknown why the IMF is universal, but it might suggest that star formation is everywhere the same.

Are there are any other clues as to the universality of star formation? Many (most?) stars form as binary and multiple systems and these provide additional information on the star formation process in a region. For example, if one region produces many wide binaries with very different masses, while another region produces fewer binaries which are all close equal-mass systems, then star formation in those two regions must have been very different, even if the IMFs are the same (Goodwin and Kouwenhoven 2009).

Observationally the binary fractions and wide binary populations of different star forming regions *are* different. For example, Taurus has a binary fraction of almost unity (Leinert et al. 1993; Patience et al. 2002), while the binary fraction of Orion is more like the field with very few wide binaries (Reipurth et al. 2007; Scally et al. 1999).

So, do different binary populations imply different star formation? The answer is maybe, but maybe not. Taurus is dynamically young and has little time to process its initial binary population so it should be close to its birth population. However, Orion is dynamically old and so the birth population may have been significantly altered (Parker et al. 2009). It is possible to model the *current* binary population of Orion as initially Taurus-like and evolved, or as initially different and less evolved (see Goodwin 2010).

Therefore the crucial question of the universality of star formation is unclear. Different regions may form stars in very different ways, or it may always be basically the same.

4 Conclusions

How stars form is an extremely complex problem. Extremely good progress has been made observationally and theoretically over the past few years, but even some quite basic issues in star formation are far from being understood.

The IMF appears universal, but its origin and why it is universal are unknown. The basics of low-mass star formation seem relatively well-understood, but initial binary fractions and distributions and the origin of binarity are unclear. High-mass star formation is very uncertain, and it is not understood how high-mass stars are able to overcome feedback to accrete to masses as high as are observed.

While the lack of understanding of many basic aspects of star formation is rather annoying, it does mean that star formation is an area that has a vast potential for new and exciting science in the coming decades.

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