THE TIMESCALE AND MODE OF STAR FORMATION IN CLUSTERS

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I discuss two questions about the origin of star clusters: How long does the process take? What is the mode of individual star formation? I argue that observations of Galactic star-forming regions, particularly the Orion Nebula Cluster (ONC), indicate that cluster formation often takes several Myr, which is many local dynamical timescales. Individual stars and binaries, including massive stars, appear to form from the collapse of gas cores.

Keywords: stars: formation — stars: pre-main sequence — open clusters and associations

1 The Timescale for Star Cluster Formation

In this section I ask: how long, in terms of dynamical timescales, does star cluster formation take? In the next, I consider how gas joins individual stars, forming in a cluster — is it from the collapse of pre-existing gas cores with masses that help determine the final stellar mass, or from the sweeping-up of gas, initially unbound to the protostar? I try to derive answers from an interpretation of observational data and then discuss the implications for theoretical models.

Star clusters are born from the densest gas clumps in giant molecular clouds (GMCs). We can measure how long this takes, from the beginning to end of star formation, in terms of the number of free-fall timescales, $t_{ff} = (3\pi/32G\rho)^{1/2}$ or the number of dynamical-crossing timescales $t_{dyn} = R/\sigma = 1.1(G\rho)^{-1/2}$, where $\sigma$ is the 1D velocity dispersion given by $\sigma^2 = \alpha_{vir} GM/(5R)$ with the observed virial parameter $\alpha_{vir} \sim 1$ (McKee & Tan 2003). In this case $t_{ff} = 0.50t_{dyn}$.

The formation timescale is an important overall constraint for theoretical models. It also determines how much dynamical relaxation occurs during formation. For $N$ equal mass stars the relaxation time is $t_{relax} \sim 0.1N/(\ln N)t_{dyn}$, i.e. about 14 crossing timescales for $N = 1000$. Using numerical experiments, Bonnell & Davies (1998) found that the mass segregation time (of clusters with mass-independent initial velocity dispersions) was similar to the relaxation time.

To evaluate the free-fall and dynamical-crossing timescales requires a choice of density and/or scale, but in reality density varies with scale. Gas clumps have approximately power-law density profiles with $\rho \propto r^{-k_\rho}$; $k_\rho \simeq 1.5-1.8$ (e.g. Mueller et al. 2002). The stellar distributions of young clusters are also usually centrally concentrated; e.g. Hillenbrand & Hartmann (1998) fit King models to the ONC. Allowing for a 50% formation efficiency, Elmegreen (2000) estimated the density in the proto-ONC as $n_H = 1.2 \times 10^5$ cm$^{-3}$, so $t_{ff} = 1.25 \times 10^5$ yr and $t_{dyn} = 2.5 \times 10^5$ yr. This applies where the present stellar density is $2 \times 10^3 M_\odot$ pc$^{-3}$, i.e. about 0.3-0.4 pc from the center. Bonnell & Davies (1998) estimated a half-mass radius of 0.5 pc, and using $\sigma = 2.5$ km s$^{-1}$ derived $t_{dyn} \sim 2 \times 10^5$ yr (note, they define a crossing time $2R/\sigma$, i.e. $4 \times 10^5$ yr).

These estimates are based on the present spatial distribution of matter, even if allowing for a formation efficiency. We expect that the initial distribution was probably more extended, since...
self-gravity and the decay of turbulent support should lead to contraction, at least in the early stages of formation. However, feedback from star formation could maintain turbulence (Norman & Silk 1980; see below). Also, dispersal of gas leads to the expansion of the cluster (Hills 1980).

To measure the actual formation time, a common method is to find the age spread of individual cluster members; for young clusters this is only practical for lower-mass stars that have a pre-main-sequence phase once they have finished accreting (high-mass stars reach the main sequence while still accreting). Observed positions in the Hertzsprung-Russell (HR) diagram are compared to theoretical models of evolution (e.g. Palla & Stahler 1999, hereafter PS99). Modeling uncertainties include the choice of deuterium abundance; D burning swells accreting protostars, so raising the “birthline” where they appear on the HR diagram (Stahler 1988). A relatively high value of \( \frac{[D/H]}{H} = 2.5 \times 10^{-5} \) was adopted by PS99. Increasing the accretion rate also raises the birthline. PS99 used constant rates of \( 10^{-5} \, M_\odot \, \text{yr}^{-1} \), which may be smaller than expected in the high pressure environments of forming clusters (McKee & Tan 2003). Finally the birthline is also influenced by the geometry of accretion (PS99 assumed spherical as opposed to disk accretion), since this affects the energetics of the gas just before it joins the stellar surface (Hartmann 2003). Fortunately, as the initial contraction from the birthline is quite rapid, the birthline position mostly affects age determinations \( < 1 \, \text{Myr} \) old. The ages of older stars tend to be estimated more robustly. Uncertainties also become larger at the lowest stellar masses.

The observables for each star are luminosity and surface temperature. Some stars are likely to be unresolved binaries, which must be allowed for (e.g. PS99). The surface temperatures and luminosities of a substantial fraction of ONC stars have been measured (Hillenbrand 1997). This sample is biased against very embedded sources and low-mass stars (\( \lesssim 0.4 M_\odot \)). Patchy extinction introduces errors. Some systematic errors are evident as the high-mass zero age main sequence lies above the theoretical expectation by 0.3 dex in luminosity (or below by 0.05 dex in temperature). If these effects also operate at lower masses, then ages would be underestimated.

PS99 could estimate ages of 258 stars with masses \( 0.4 < m_\ast / M_\odot < 6.0 \) from Hillenbrand’s (1997) sample. The lower limit of this range was set because of incompleteness at \( L_\ast \simeq 0.1 L_\odot \), the predicted luminosity of a 10 Myr old \( 0.4 M_\odot \) star. However, given that the model uncertainties are largest at these low masses, and given possible observational systematic uncertainties, the sample may not be complete at \( 0.4 M_\odot \) to ages of only several Myr. The large numbers of stars at these low masses increases the potential importance of incompleteness in the lowest mass bin. When divided into different mass intervals, the lowest mass bin with mean mass of \( \bar{m}_\ast = 0.56 M_\odot \), does appear to be different from the next two bins with \( \bar{m}_\ast = 1.0, 1.8 M_\odot \).

Assuming a complete sample, PS99 found: 82 stars aged 0-1 Myr, 57 aged 1-2 Myr, 34 aged 2-3 Myr, 17 aged 3-4 Myr, 8 aged 4-5 Myr, 8 aged 5-6 Myr, 8 aged 6-7 Myr, and 6 aged 7-10 Myr. Hartmann (2003) has argued that the oldest ages (\( \sim 10 \, \text{Myr} \)) may be due to a problem of foreground contamination. As described above, birthline uncertainties mostly affect the youngest ages (\( \lesssim 1 \, \text{Myr} \)). We conclude that a significant fraction of the stars are as old as 3 Myr, and that this is a lower limit to \( t_{\text{form}} \) since star formation is still continuing in the cluster.

An independent method of dating the cluster comes from the identification of a dynamical ejection event of 4 massive stars (a binary and two singles) that appear to have originated from the ONC about 2.5 Myr ago (Hoogerwerf, de Bruijne & de Zeeuw 2001). The central value of the time since this ejection event is about 2.3 Myr in the analysis of Hoogerwerf et al. (2001); however, if the cluster’s distance of about 450 pc is adopted, then the best estimate for the age is 2.5 Myr. The identification is based upon the fact that the extrapolation of the motion of the center of mass of the four stars from the ejection event to the present day, leads to a predicted position coincident with the ONC (uncertainties are a couple of pc). These results imply that 2.5 Myr ago the ONC was already a rich cluster containing at least four stars of spectral type earlier than O9/B0. Before this the stars had to form and have enough time to find and eject each other in a close interaction. Thus the estimate of 2.5 Myr is again a lower limit to \( t_{\text{form}} \).
Taken together, the above results lead us to conclude, with some confidence, that the formation timescale for the cluster is $t_{\text{form}} \geq 3$ Myr. This is $\geq 12$ (24) dynamical-crossing (free-fall) timescales, as estimated at the conditions considered by Elmegreen (2000) (see above). This age corresponds to a free-fall timescale for densities of $\rho = 210$ cm$^{-3}$, much smaller than the mean density of the ONC region. The dynamical ejection event implies that a dense stellar cluster existed 2.5 Myr ago, i.e. that the densities have not changed appreciably during this time, and the ONC has taken many ($\geq 10$) dynamical-crossing times to form.

How does this result compare to what is known from other star-forming regions? Forbes (1996) did not find evidence for an age spread in NGC 6531, but the analysis was insensitive to timescales shorter than at least 3 Myr. Hodapp & Deane (1993) analyzed L1641, placing twelve stars in an HR diagram: ten are spread in age from 0-2 Myr and two are about 6 Myr old. The conclusions that can be drawn from this cluster are limited due to the small sample, lack of correction for binarity, and use of relatively old pre-main-sequence tracks. Palla & Stahler (1997) and Belikov et al. (1998) concluded that the age spread in the Pleiades was $\sim$ 30 Myr around a mean value of about 100 Myr. If true, then this is the largest measured formation time of local clusters.

If the formation time is many dynamical timescales, then there should be virialized and heavily embedded clusters at early stages in their evolution (most clusters analyzed with the HR diagram method are by necessity optically revealed and probably nearing the end of star formation). This population of embedded clusters probably corresponds to the population of hot molecular clumps observed in sub-mm continuum and molecular lines (e.g. Mueller et al. 2002; Shirley et al. 2003). Indeed most of the CS line maps of Shirley et al. (2003) have morphologies consistent with quasi-spherical virialized distributions. The clumps have typical masses $M \sim 100 - 10^4$ $M_\odot$, diameters $\sim 1$ pc and surface densities $\Sigma \sim 1$ g cm$^{-2}$.

Relatively slow, almost quasi-static evolution of the star-forming clump, contrasts with some theories and models of star formation that take only one or a few crossing times (e.g. Elmegreen 2000; Bonnell & Bate 2002). One theoretical motivation for such fast timescales has been the rapid decay of turbulence: even in a strongly magnetized medium, the kinetic energy associated with supersonic turbulence decays with a half-life of just over a signal crossing timescale (Stone, Ostriker, & Gammie 1998). Thus a long formation time requires that the turbulence observed in star-forming clumps is maintained by energy input, most likely from protostellar outflows and the overall contraction of the clump. In fact the energy requirements are quite modest; the energy dissipation rate of virialized clumps (with $k_\rho = 3/2$ and mean surface density $\Sigma = M/(\pi R^2)$) that lose half their kinetic energy in one dynamical crossing time, $t_{\text{dyn}}$ is $21(M/4000 M_\odot)^{5/4} \Sigma^{5/4} L_\odot$. Even very inefficient coupling of protostellar outflows to the gas allows turbulence to be maintained. The initial kinetic energy of outflows is expected to be about half of the total energy released associated with accretion (e.g. Shu et al. 2000), but much of this is dissipated in shocks as gas is swept-up. The collimated nature of the flows also means that a lot of energy escapes from the star-forming region. Nevertheless, even if only 1% of the outflow energy generates turbulence in the clump, then the cluster can form leisurely in 20 dynamical-crossing times and still maintain its turbulent support.\footnote{This calculation assumes 50% star formation efficiency, a turbulent decay half-life of 1 $t_{\text{dyn}}$, a Salpeter initial mass function (IMF) from 0.1 to 120 $M_\odot$, protostellar sizes based on the model of McKee & Tan (2003), outflow mechanical luminosity of $Gm_\star m_\star/(2r_\star)$, and a coupling efficiency of 0.01. The formation time that maintains turbulent support is then $t_{\text{form}} = \eta_{\text{dyn}} t_{\text{dyn}}$ with $\eta_{\text{dyn}} = 21(M/4000 M_\odot)^{-1/2} \Sigma^{-1/2}$.} This inefficiency of star
formation is consistent with some numerical simulations of self-gravitating, unmagnetized gas with driven turbulence (Vázquez-Semadeni, Ballesteros-Paredes, & Klessen 2003).

Note that while the observed age spreads of clusters imply formation times long compared to the local dynamical timescale of the clump, these times are about equal to the dynamical timescale of the larger scale GMCs, in which clumps are embedded. Hartmann, Ballesteros-Paredes, & Bergin (2001) used this fact to argue that GMCs are transient objects, which would be true if GMCs did not have a long period of quiescence before star formation and if they were destroyed quickly once the stars had formed. However, the star formation in most GMCs is spatially localized, so that most of the mass is quiescent, and it is not clear that most newly formed star clusters are able to destroy their parental clouds. The observational study of Leisawitz, Bash & Thaddeus (1989) found that open star clusters older than about $\sim 10$ Myr were not associated with molecular clouds, which is consistent either with post-star-formation cloud lifetimes shorter than this age or with relative velocities of star clusters and their parent clouds of about $10 \text{ km s}^{-1}$, as might arise from photoionization feedback (Williams & McKee 1997).

One traditional objection to longer formation timescales has been the idea that once massive stars are formed in a cluster, they would very rapidly disperse the gas with their ionizing radiation and stellar winds. Observational evidence suggests that massive stars are not always the last stars to form in their clusters (e.g. Hoogerwerf et al. 2001). Tan & McKee (2001; 2004) found that the turbulent, clumpy, and self-gravitating nature of gas led to much slower dispersal times, compared to a uniform medium. A single O star producing $10^{49}$ H-ionizing photons s$^{-1}$ was unable to disperse the gas in a 4000 $M_\odot$ clump with a density comparable to the proto-ONC. If the gas in the clump was allowed to form stars at a rate such that 50% of the initial mass would be turned into stars in $15 t_{\text{dyn}}$ and the feedback from this star formation was accounted for, then the gas was dispersed in about 2 Myr (about $10 t_{\text{dyn}}$). The dynamical ejection of massive stars from clusters, which was not included in the above calculations, would increase the dispersal timescale. We have seen that such an ejection occurred in the ONC 2.5 Myr ago. It also appears to be happening at the present epoch with the ejection of the Becklin-Neugebauer (BN) object and $\Theta^1 C$, the most massive star in the cluster (Plambeck et al. 1995; Tan 2004a).

A 3 Myr formation time for the ONC has implications for how much mass segregation has occurred, which bears upon the question of whether massive stars tend to form preferentially in the centers of clusters. This time corresponds to about 8 diameter crossing times at the half mass radius of 0.5 pc, which is the unit of time used in the study of Bonnell & Davies (1998). If the 30 most massive stars are initially placed at the half-mass radius, $r_{1/2}$, then after this time the median location of the 6 most massive stars has migrated in to only $0.15r_{1/2}$. However, the effects of gradual formation and the presence of gas need to be accounted for before detailed comparison is made to the present-day ONC stellar distribution.

2 The Mode of Star Formation in Star Clusters

The paradigm of star formation from initially quasi-hydrostatic gas cores (e.g. Shu, Adams, & Lizano 1987) has faced the challenges of competitive accretion and dynamical interactions when applied to star formation, particularly of high-mass stars, in nascent star clusters (e.g. Bonnell, Bate, & Zinnecker 1998; Stahler, Palla, & Ho 2000). For example, the smooth-particle-hydrodynamics (SPH) simulations of Bonnell, Bate and collaborators are characterized by a much more chaotic evolution in which long-lived cores are not readily apparent. Bonnell, Vine, & Bate 2004 showed that the most massive star at the end of their simulation had gained mass that was initially very widely distributed. However, cores (both starless and with embedded protostars) are observed in real proto-clusters (e.g. Testi & Sargent 1998; Motte et al. 2001; Li, Goldsmith, & Menten 2003; Beuther & Schilke 2004), and even have an IMF that, although uncertain, is similar to that of stars. Part of the resolution of this discrepancy may lie in the
methodology of the numerical simulations: these do not yet include magnetic fields or feedback from forming stars; the equation of state is isothermal; the SPH method uses a softening length for each particle so does not resolve features smaller than this scale, particularly shocks; and protostars are modeled as sink particles that form when a mass of $\geq 0.1 M_\odot$ happens to be self-gravitating, sub-virial and at a density $n_{H_2} \geq 3.1 \times 10^8$ cm$^{-3}$. One example of where the SPH technique may fail is the problem of the rate of Bondi-Hoyle accretion, which is how most of the stellar mass is acquired in the above simulations. In this process gas is gravitationally focused by a passing star so that streamlines collide, shock and dissipate their energy. It occurs on scales unresolved by the present SPH techniques. Eulerian grid simulations, including adaptive mesh refinement of small scale structures, have been used to simulate the interaction of sink particles with surrounding turbulent gas: the accretion rate is much smaller than the classical analytic estimate of accretion from a uniform medium (Krumholz, McKee, & Klein 2004).

In spite of recent progress in the observations of high-mass star-forming regions, the mode of massive star formation remains controversial. There is as yet no clear-cut example of a massive protostar for which the accretion disk and core are readily identified. One difficulty is the uncertainty in the expected properties of such disks and cores. McKee & Tan (2003) described the properties of virialized cores in pressure equilibrium with their high-pressure proto-cluster surroundings, and calculated the evolution of the protostars that form from them. Tan (2004b) compared these models to observations of the Orion Hot Core in the Kleinmann-Low region of the ONC, concluding that the picture of massive star formation from the collapse of massive cores was supported. Here we highlight the main points of this argument.

The total luminosity of the Kleinmann-Low region is about $10^5 L_\odot$, and the orientation of the polarization vectors of 3.8 $\mu$m emission suggest that much of this comes from a very localized region close to the center of a dense and hot molecular core, known as the Orion Hot Core (Werner, Dinnerstein, & Capps 1983). Thermal emission from dust in this structure is also seen at 450 $\mu$m and it has a radius of $\sim 0.05$ pc (Wright et al. 1992), about that expected for an initially 60 $M_\odot$ core. If about half of the total luminosity of the region comes from a single protostar, then it would have a mass of about 20 $M_\odot$ (McKee & Tan 2003), so, allowing for accretion inefficiencies due to outflows, the core is about half way through its collapse. If the initial ratio of rotational to gravitational energy is about 2%, as is typical for low-mass cores (Goodman et al. 1993), then the expected size of the accretion disk is about 1000 AU. This corresponds to the size of the region of SiO (v=0) maser emission seen by Wright et al. (1995). If the velocity gradient across this structure is interpreted as being due to Keplerian motion of a disk, then the central mass is about 30 $M_\odot$. The orientation of the disk is perpendicular to the larger scale molecular outflow that is ejected to the NW and SE of this region (e.g. Chernin & Wright 1996). At the center of the SiO (v=0) emission is a thermal radio source, known as “I” (e.g. Menten & Reid 1995). It is elongated perpendicular to the disk (Menten & Reid, in preparation), i.e. parallel to the outflow. In particular, the major axis aligns very closely with the direction towards Herbig-Haro objects to the NW that show blue-shifted velocities of up to 400 km s$^{-1}$ (Taylor et al. 1986). Allowing for flow geometry, this corresponds to a total velocity of 1000 km s$^{-1}$, which is the expected initial outflow velocity from a $\sim 20 - 30 M_\odot$ protostar, i.e. about its escape speed. The thermal radio source is likely due to ionized gas: the massive protostar is only able to ionize a small patch of its outflow, which is dense enough to confine the radiation, except along the rotation axis. These “outflow-confined” HII regions have been considered by Tan & McKee (2003): the models explain the elongation and radio spectrum of source “I”. In summary, there are many independent pieces of evidence that corroborate a model of star formation involving collapse of a massive core to an accretion disk, which then feeds a massive ($\sim 20 M_\odot$) protostar, driving a powerful outflow in the process. No individual piece of evidence is conclusive, but in their totality a remarkably consistent picture emerges.

The situation is, of course, somewhat more complicated than the above description. There
are a few other protostars and outflows in the region, but these seem to have quite modest luminosities and their apparent proximity to the Orion Hot Core may be due to projection effects. If there is a physically close second protostar, then this may indicate that the core is collapsing to form a binary, a relatively minor extension of the basic model. In addition, the collapse of the core is probably perturbed by other cluster members; e.g., the BN object made a close projected passage to source “I” about 500 yr ago (Tan 2004a). On smaller scales of several tens of AU from the center of source “I”, SiO (v=1) maser emission is seen with a morphology approximately in the shape of an X, stretched along the NW-SE axis (Greenhill et al. 2003). The densities and temperatures required for excitation are similar to those expected in the inner part of the outflow (Tan & McKee 2003), but the velocity structure has also been interpreted as evidence for a disk aligned along this NW-SE axis (Greenhill et al. 2003), which would be incompatible with the above model. The small velocity differences (∼ tens km s\(^{-1}\)) in the maser spots are hard to understand in the context of an outflow from a massive star that should have much faster speeds, although the radial velocity range surveyed so far is only ∼ 100 km s\(^{-1}\). Further observations of these features would be very useful to help resolve this issue.

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