Star Formation

Fig. 3.—Spherical slice of the gas density inside the Jeans volume at \( \tau = 1 \) with 128 cells per Jeans length \( y \) at velocity streamlines on a linear color scale ranging from dark blue to light gray. Magnetic field lines showing a highly tangled and twisted magnetic field structure typical of the small-scale dynamo; yellow: 0.5 m G, red: 1 m G. Four randomly chosen individual field lines: The green one in particular is extremely tangled close to the center of the Jeans volume. Contours of the vorticity modulus \( |\vec{\omega}| \) show elongated filamentary structure typical for subsonic turbulence. Frisch (1995) for spherical slice of the divergence of the velocity field \( \nabla \cdot \vec{v} \); white: compression, red: expansion.
agenda

• star formation theory
  - phenomenology
  - historic remarks
  - our current understanding and its limitations

• application
  - the stellar mass function at birth (IMF)
phenomenology
• star formation sets in very early after the big bang
• stars always form in galaxies and protogalaxies
• we cannot see the first generation of stars, but maybe the second one
M51 with Hubble (additional processing R. Gendler)
• correlation between stellar birth and large-scale dynamics
• spiral arms
• tidal perturbation from neighboring galaxy
galaxies from THINGS and HERACLES survey
(images from Frank Bigiel, ZAH/ITA)
• HI gas more extended
• H2 and SF well correlated

galaxies from THINGS and HERACLES survey (images from Frank Bigiel, ZAH/ITA)
distribution of molecular gas in the Milky Way as traced by CO emission
data from T. Dame (CfA Harvard)

Orion Nebula Cluster (ESO, VLT, M. McCaughrean)
Orion Nebula Cluster (ESO, VLT, M. McCaughrean)
• stars form in molecular clouds
• stars form in clusters
• stars form on ~ dynamical time
• (protostellar) feedback is very important
Ionizing radiation from central star $\Theta$ 1C Orionis Trapezium stars in the center of the ONC (HST, Johnstone et al. 1998)
• strong feedback: UV radiation from \( \Theta 1 \)C Orionis affects star formation on all cluster scales

Trapezium stars in the center of the ONC (HST, Johnstone et al. 1998)
eventually, clusters like the ONC (1 Myr) will evolve into clusters like the Pleiades (100 Myr)
theoretical approach
• density
  - density of ISM: few particles per cm\(^3\)
  - density of molecular cloud: few 100 particles per cm\(^3\)
  - density of Sun: 1.4 g/cm\(^3\)

• spatial scale
  - size of molecular cloud: few 10s of pc
  - size of young cluster: \(\sim 1\) pc
  - size of Sun: \(1.4 \times 10^{10}\) cm
• contracting force
  - only force that can do this compression is GRAVITY

• opposing forces
  - there are several processes that can oppose gravity
    - GAS PRESSURE
    - TURBULENCE
    - MAGNETIC FIELDS
    - RADIATION PRESSURE

Modern star formation theory is based on the complex interplay between all these processes.
early theoretical models

● *Jeans (1902):* Interplay between self-gravity and thermal pressure

- stability of homogeneous spherical density enhancements against gravitational collapse
- dispersion relation:

\[ \omega^2 = cs^2 k^2 - 4\pi G \rho_0 \]

- instability when \( \omega^2 < 0 \)
- minimal mass:

\[ M_J = \frac{1}{6} \pi^{-5/2} G^{-3/2} \rho_0^{-1/2} c_s^3 \propto \rho_0^{-1/2} T^{3/2} \]
first approach to turbulence

- von Weizsäcker (1943, 1951) and Chandrasekhar (1951): concept of **MICROTURBULENCE**
  
  - BASIC ASSUMPTION: separation of scales between dynamics and turbulence
    \[ l_{\text{turb}} \ll l_{\text{dyn}} \]
  
  - then turbulent velocity dispersion contributes to effective soundspeed:
    
    \[ \frac{C_c^2}{\sigma_{\text{rms}}} \rightarrow \frac{C_c^2}{\sigma_{\text{rms}}} + \frac{\sigma_{\text{rms}}^2}{\text{rms}} \]

  - → Larger effective Jeans masses → more stability
  
  - BUT: (1) turbulence depends on \( k \): \( \sigma_{\text{rms}}^2(k) \)
  
  (2) supersonic turbulence \( \rightarrow \sigma_{\text{rms}}^2(k) \gg \omega_0^2 \)
problems of early dynamical theory

● molecular clouds are *highly Jeans-unstable*, yet, they do *NOT* form stars at high rate and with high efficiency (Zuckerman & Evans 1974 conundrum) (the observed global SFE in molecular clouds is $\sim 5\%$) \rightarrow *something prevents large-scale collapse.*

● all throughout the early 1990’s, molecular clouds had been thought to be long-lived quasi-equilibrium entities.

● molecular clouds are *magnetized*
Magnetic star formation

Mestel & Spitzer (1956): Magnetic fields can prevent collapse!!!

- Critical mass for gravitational collapse in presence of B-field

\[
M_{cr} = \frac{5^{3/2}}{48\pi^2} \frac{B^3}{G^{3/2} \rho^2}
\]

- Critical mass-to-flux ratio
  (Mouschovias & Spitzer 1976)

\[
\left[ \frac{M}{\Phi} \right]_{cr} = \frac{\xi}{3\pi} \left[ \frac{5}{G} \right]^{1/2}
\]

- Ambipolar diffusion can initiate collapse

Lyman Spitzer, Jr., 1914 - 1997
“standard theory” of star formation

• BASIC ASSUMPTION: Stars form from magnetically highly subcritical cores

• Ambipolar diffusion slowly increases \((M/\Phi)\): \(\tau_{AD} \approx 10\tau_{ff}\)

• Once \((M/\Phi) > (M/\Phi)_{\text{crit}}\):
  dynamical collapse of SIS
  • Shu (1977) collapse solution
  • \(dM/dt = 0.975 c_s^3/G = \text{const.}\)

• Was (in principle) only intended for isolated, low-mass stars
problems of “standard theory”

- Observed B-fields are weak, at most marginally critical (Crutcher 1999, Bourke et al. 2001)


- Structure of prestellar cores (e.g. Bacman et al. 2000, Alves et al. 2001)

- Strongly time varying dM/dt (e.g. Hendriksen et al. 1997, André et al. 2000)

- More extended infall motions than predicted by the standard model (Williams & Myers 2000, Myers et al. 2000)

- Most stars form as binaries (e.g. Lada 2006)

- As many prestellar cores as protostellar cores in SF regions (e.g. André et al 2002)

- Molecular cloud clumps are chemically young (Bergin & Langer 1997, Pratap et al 1997, Aikawa et al 2001)

- Stellar age distribution small ($\tau_{ff} \ll \tau_{AD}$) (Ballesteros-Paredes et al. 1999, Elmegreen 2000, Hartmann 2001)

- Strong theoretical criticism of the SIS as starting condition for gravitational collapse (e.g. Whitworth et al 1996, Nakano 1998, as summarized in Klessen & Mac Low 2004)

- Standard AD-dominated theory is incompatible with observations (Crutcher et al. 2009, 2010ab, Bertram et al. 2011)

(see e.g. Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)
BASIC ASSUMPTION:
star formation is controlled by interplay between supersonic turbulence and self-gravity

turbulence plays a dual role:
- on large scales it provides support
- on small scales it can trigger collapse

some predictions:
- dynamical star formation timescale $\tau_{ff}$
- high binary fraction
- complex spatial structure of embedded star clusters
- and many more . . .

Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194
McKee & Ostriker, 2007, ARAA, 45, 565
molecular clouds
\[ \sigma_{\text{rms}} \approx \text{several km/s} \]
\[ M_{\text{rms}} > 10 \]
\[ L > 10 \text{ pc} \]

energy source & scale
\textit{NOT known}
(supernovae, winds, spiral density waves?)

massive cloud cores
\[ \sigma_{\text{rms}} \approx \text{few km/s} \]
\[ M_{\text{rms}} \approx 5 \]
\[ L \approx 1 \text{ pc} \]

dense protostellar cores
\[ \sigma_{\text{rms}} \ll 1 \text{ km/s} \]
\[ M_{\text{rms}} \leq 1 \]
\[ L \approx 0.1 \text{ pc} \]

dissipation scale not known
(ambipolar diffusion, molecular diffusion?)

turbulent cascade in the ISM
turbulence creates a hierarchy of clumps
as turbulence decays locally, contraction sets in
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while region contracts, individual clumps collapse to form stars
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individual clumps collapse to form stars
individual clumps collapse to form stars
in dense clusters, clumps may merge while collapsing
--> then contain multiple protostars
in *dense clusters*, clumps may merge while collapsing
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--> then contain multiple protostars
in *dense clusters*, competitive mass growth becomes important
in *dense clusters*, competitive mass growth becomes important
in dense clusters, $N$-body effects influence mass growth
low-mass objects may become ejected --> accretion stops
feedback terminates star formation
result: *star cluster*, possibly with H\textsubscript{II} region
current status

- stars form from the complex interplay of self-gravity and a large number of competing processes (such as turbulence, B-field, feedback, thermal pressure)

- the relative importance of these processes depends on the environment
  - prestellar cores --> thermal pressure is important
  - molecular clouds --> turbulence dominates

- massive star forming regions (NGC602): radiative feedback is important
- small clusters (Taurus): evolution maybe dominated by external turbulence

- star formation is regulated by various feedback processes
- star formation is closely linked to global galactic dynamics (KS relation)

Star formation is intrinsically a multi-scale and multi-physics problem, where it is difficult to single out individual processes. Simple theoretical approaches usually fail.
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selected open questions

- what processes determine the initial mass function (IMF) of stars?
- what are the initial conditions for star cluster formation? how does cloud structure translate into cluster structure?
- how do molecular clouds form and evolve?
- what drives turbulence?
- what triggers / regulates star formation on galactic scales?
- how does star formation depend on metallicity? how do the first stars form?
- star formation in extreme environments (galactic center, starburst, etc.), how does it differ from a more “normal” mode?
stellar mass function
stellar mass function

stars seem to follow a universal mass function at birth --> IMF

Orion, NGC 3603, 30 Doradus
(Zinnecker & Yorke 2007)
stellar masses

- distribution of stellar masses depends on
  - turbulent initial conditions
    --> mass spectrum of prestellar cloud cores
  - collapse and interaction of prestellar cores
    --> accretion and $N$-body effects
  - thermodynamic properties of gas
    --> balance between heating and cooling
    --> EOS (determines which cores go into collapse)
  - (proto) stellar feedback terminates star formation
    ionizing radiation, bipolar outflows, winds, SN

(Kroupa 2002)
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(Kroupa 2002)
example: model of Orion cloud

„model“ of Orion cloud:
15.000.000 SPH particles,
$10^4 \, M_{\text{sun}}$ in 10 pc, mass resolution
$0,02 \, M_{\text{sun}}$, forms $\sim 2.500$
„stars“ (sink particles)

isothermal EOS, top bound, bottom unbound

has clustered as well as distributed „star“ formation

efficiency varies from 1% to 20%

develops full IMF
(distribution of sink particle masses)

(Bonnell, Smith, Clark, & Bate 2010, MNRAS, 410, 2339)
example: model of Orion cloud

Bonnell et al. 2010

example: model of Orion cloud

(Massive stars) form early in high-density gas clumps (cluster center) - high accretion rates, maintained for a long time

Low-mass stars - form later as gas falls into potential well - high relative velocities - little subsequent accretion
• distribution of stellar masses depends on
  - turbulent initial conditions
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application to early star formation
degree of fragmentation depends on EOS!

polytropic EOS: $p \propto \rho^\gamma$

$\gamma < 1$: dense cluster of low-mass stars

$\gamma > 1$: isolated high-mass stars

(see Li et al. 2003; also Kawachi & Hanawa 1998, Larson 2003)
dependency on EOS

for $\gamma < 1$ fragmentation is enhanced $\Rightarrow$ cluster of low-mass stars

for $\gamma > 1$ it is suppressed $\Rightarrow$ isolated massive stars

(from Li, Klessen, & Mac Low 2003, ApJ, 592, 975)
how does that work?

(1) \( p \propto \rho^\gamma \rightarrow \rho \propto p^{1/\gamma} \)

(2) \( M_{\text{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2} \)

- \( \gamma < 1 \): \( \rightarrow \) large density excursion for given pressure
  \( \rightarrow \) \( \langle M_{\text{jeans}} \rangle \) becomes small
  \( \rightarrow \) number of fluctuations with \( M > M_{\text{jeans}} \) is large

- \( \gamma > 1 \): \( \rightarrow \) small density excursion for given pressure
  \( \rightarrow \) \( \langle M_{\text{jeans}} \rangle \) is large
  \( \rightarrow \) only few and massive clumps exceed \( M_{\text{jeans}} \)
EOS as function of metallicity

In this section, we review thermal evolution of the cloud core in different metallicities. The cooling and heating rates by individual processes are compared for different metallicities. This should be connected with the metallicity effects on the dynamical collapse.

For the cooling, the H$_2$ line emission contributes to the temperature evolution at each metallicity. The contribution of specific heat at the center, one-zone models are presented. In Figure 2 of O05, where similar plots for the overall evolution is quite similar to that calculated by the one-zone model (Figure 3.2), justifying the one-zone treatment for cases with metals. Below $10^{-5}$, metallicity effects are so small that the temperature evolution is almost identical to the metal-free one except for a slight offset at highest densities.

The heating is owing to the compression of the gas. With gradual increase of temperature, the balance of chemical reaction via the three-body H$_2^+ + H ightarrow H_2 + e$ channel is quenched. For the cooling, the H$_2$ collision-induced continuum absorption. Another molecular species in the metal-free case is the H$_2^+$ channel by the density catalyzed by a small amount of remaining electrons. With increasing density, the hydrogen is converted to the molecular form via this channel. After this plateau, the H$_2$ line emission contributes comparably to H$_2$ and this suppresses the cooling rate.

The temperature evolution at the center of cloud cores during the core evolution. There are, however, small disagreements, in particular, at high densities and for low-metallicity cases. Note that the temperature $\sim 10^3$ K is shown for those cases. For example, the clouds easily fragment as long as this value, the dynamical collapse is halted as the pressure becomes so small that the temperature evolution is almost identical to the metal-free one except for a slight offset at highest densities.

The evolution of H$_2$ number density log ($n_H$ cm$^{-3}$) for different metallicities. This should be connected with the metallicity effects on the dynamical collapse.

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Figure 1. Gas is assumed to be fully atomic (molecular) in drawing those lines. Compared with Figure 2 of O05, where similar plots for the zone model (Figure 11) present the temperature evolution at each metallicity. The contribution to the cooling and heating rates by individual processes are small that the evolution is almost identical to the metal-free one except for a slight offset at highest densities (Omukai et al. 2005, 2010).

In this section, we review thermal evolution of the cloud core over time. In the case of metallicity $[M/H] < -5$, metallicity effects are so small that the temperature evolution is almost identical to the metal-free one except for a slight offset at highest densities. Another critical value is $\tau = 10^{-4}M_{\odot}$, which marks the fragmentation limit for high densities and for low-metallicity cases.

In particular, at high densities and for low-metallicity cases, the temperature $T(K)$ is shown for those cases. Note that $\gamma$ is the effective ratio of specific heat, which gives $\ln H/\ln T$. We then describe the effects of metallicity on these differences later in Section 3.2.
Evolution of temperatures at the center of cloud cores during the prestellar collapse for various metallicities. This is calculated by one-zone models. The contribution of the curve in Figure 10 compared to the cooling and heating rates by individual processes are presented in Figure 3.4. We defer detailed discussion on these differences to later in the text.

Let us summarize here the formation processes of H\(_2\) in metal-free gas. We then describe the effects of metallicity on the dynamical response of self-gravitating clouds to thermal heating and cooling.

Next, let us see the cooling and heating processes associated with the three-body reaction (Equation (11)) and its equilibrium between the H\(_2\) and H\(_3^+\) channels:

\[ \text{H}_2 + 	ext{H} \rightarrow \text{H}_3^+ + e^- \]

After this plateau, the H\(_2\) formation (Equation (10)) starts to increase, mainly catalyzed by a small amount of remaining electrons. With gradual increase of temperature, the balance of chemical reactions becomes important, and a plateau is reached. The rate of H\(_2\) formation increases with temperature until it saturates at a brief period. The rise in the formation heating rate at 10\(^{-16}\) cm\(^{-3}\) is due to spatially limited reactions between H\(_2\) and H\(_3^+\) (Equation (12)).

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EOS as function of metallicity

For those above 10^8 M_sun, the dynamical response of self-gravitating clouds to thermal evolution. The effective ratio of specific heat is an important index to examine the core evolution. There are, however, small disagreements, (Omukai et al. 2005, 2010) the metal-free one except for a slight offset at highest densities are so small that the temperature evolution is almost identical to the metal-free case. In the case of metallicity [M/H]=-6, HD, is known to play an important role in cooling associated with the three-body reaction (Equation (12)).

Let us summarize here the formation processes of H_2, justifying the one-zone treatment for the cloud. After this plateau, the H_2 abundance begins to increase again at densities where the H_2 channel is quenched (O05, where similar plots for the cases with metals. Below 10^8 M_sun, fragmentation is strongly prohibited for the metal-free gas, HD, is known to play an important role in cooling.

Two critical values are [M/H]=-6 and [M/H]=-3, but the thermal process control is due to the H_2 channel by the density.

The effective ratio of specific heat at the center, compared with Figure 2 of O05, where similar plots for the金属 cases are presented in Figure 3. A color version of this figure is available in the online journal.
present-day star formation

\( Z = 0 \)

\( \tau = 1 \)

(Larson 1985, Larson 2005)

\( \gamma = 1.1 \)

\( \gamma = 0.7 \)

(Larson 1985, Larson 2005)
with $\rho_{\text{crit}} \approx 2.5 \times 10^5 \text{ cm}^{-3}$

at SFE $\approx 50$

$M_{\text{low}} / M_{\text{tot}} = 45\%$

need appropriate EOS in order to get low mass IMF right

EOS as function of metallicity

![Graph showing EOS as function of metallicity](image)

(Omukai et al. 2005, 2010)
The EOS as a function of metallicity indicates the constant Jeans masses. For those above \(10^{-4} \text{M}_{\odot}\), the gas is assumed to be fully atomic (molecular) in drawing those lines. The dashed lines in Figure 1 are so small that the temperature evolution is almost identical to the case of metallicity \([M/\text{H}]=-6\) (Omukai et al. 2005, 2010). Another critical value is \(\sim 10^{-2} \text{M}_{\odot}\), along with its isothermal line cooling formation (Equation (10)). Until very high density is reached, cooling and heating are always almost balanced, so the evolution is nearly isothermal with temperature differences below unity (Figure 2; McGreer & Bryan 2007). Although some lines become optically thick at a brief period at \(10^{-3} \text{M}_{\odot}\), the \(H\) line formation dominates. For the cooling, the \(H\) line channel is quenched if a metal-free gas is once ionized (Uehara & Inutsuka 2008), while fragmentation is strongly prohibited for \(Z/Z_{\odot}>10\). For \([M/\text{H}]=-5\), the effective ratio of specific heat is an important index to examine the dynamical response of self-gravitating clouds to thermal processes, particularly at high densities and for low-metallicity cases. The core evolution. There are, however, small disagreements, in particular, at high densities and for low-metallicity cases. We defer detailed discussion on these differences to later in this section. We review thermal evolution of the cloud core as a function of the number density. The variation of pressure in response to the density variation, and this suppresses the cooling rate gradient. The steep decline of the \(H\) line contributes comparably to \(H\) line cooling, but for \(10^{-3} \text{M}_{\odot}\), only a small factor whereas density increases by many orders of magnitude. All the hydrogen is converted to the molecular form via this channel by the density gradient that the evolution is nearly isothermal with temperature differences below unity (Figure 2; McGreer & Bryan 2007).
transition: Pop III to Pop II.5

two competing models:

- cooling due to atomic fine-structure lines \((Z > 10^{-3.5} \, Z_{\odot})\)
- cooling due to coupling between gas and dust \((Z > 10^{-5} \ldots -6 \, Z_{\odot})\)

- which one explains origin of extremely metal-poor stars?

NB: lines would only make very massive stars, with \(M > \text{few} \times 10 \, M_{\odot}\).

(Omukai et al. 2005, 2010)
SDSS J1029151+172927

- is first ultra metal-poor star with $Z \sim 10^{-4.5} Z_{\odot}$ for all metals seen (Fe, C, N, etc.)
  [see Caffau et al. 2011]
- this is in regime, where metal-lines cannot provide cooling
  [e.g. Schneider et al. 2011, 2012, Klessen et al. 2012]

- new ESO large program to find more of these stars (120h x-shooter, 30h UVES)
  [PI E. Caffau]
modeling the formation of the first/second stares

successive zoom-in calculation from cosmological initial conditions (using SPH and new grid-code AREPO)


Dependence of gas and dust temperatures on gas density for metallicities $\text{[M/H]} = -4$ to $\text{[M/H]} = -6$.

Fragmentation of star-forming clouds at very low metallicities $\text{[M/H]} =$ -4 to -6. We have performed a set of four simulations for different metallicities in order to test if dust can efficiently cool the gas and change the fragmentation behavior. Since dust cooling is more efficient for higher densities, we expect that its consequence of inelastic gas-grain collisions, and these collisions increases the number of Jeans masses present in the collapsing region, making the gas unstable to fragmentation. The values were calculated just before the formation of the first sink particle.

In the bottom panel of Figure 10, we show the accretion properties for the new mass resolution, which is 10 times higher than the mass resolution used in the previous simulation. To resolve the fragmentation, the mass resolution is increased to roughly 400 K in the $10^{-5} \, \text{Z}_\odot$ case. This temperature drop significantly increases to a few thousand Kelvin, preventing the gas temperature from getting higher than 1500 K. For instance, the metal-free case reaches temperatures close to 6000 K, while for $Z_{\odot}$, dust cooling begins to be important for densities as high as $10^{-4} \, \text{Z}_\odot$, the density where dust cooling begins to become important. And for densities as high as $10^{-3} \, \text{Z}_\odot$, dust cooling becomes important for the evolution is close to isothermal. Changes in metallicity influence the evolution of the dust and gas temperatures in the simulations, as expressed in the temperature-density diagram (Figure 1). In order to explain these differences, we take a closer look at the cooling and heating processes. While for $Z_{\odot}$, heating dominates, the evolution is close to adiabatic. And for densities as high as $10^{-3} \, \text{Z}_\odot$, dust cooling becomes important to explain the cooling processes.

The gas thermal evolution during the collapse takes different paths depending on the metallicity, as expressed in the temperature-density diagram (Figure 1). In all cases, such as for the solar value, calculated just before the first sink particle was born stellar systems. The top panel shows how the total mass luminosity. We did not take this thermal process into account creating more sparse overdensities. In Figure 10 we present accretion properties for the new mass resolution. To resolve the fragmentation, the mass resolution is increased to roughly 400 K in the collapsing region, making the gas unstable to fragmentation. The values were calculated just before the formation of the first sink particle.

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Hints for differences in mass spectrum

Disk fragmentation mode

Gravoturbulent fragmentation mode

Fig. 6.—Sink particle mass function at the point when 4.7 $M_\odot$ of gas had been accreted by the sink particles in each simulation. To resolve the fragmentation, the mass resolution is smaller than the Jeans mass at the point in the temperature-density diagram where dust and gas couple and the compressional heating starts to dominate over the dust cooling.

Fig. 7.—Timescales for fragmentation (bottom panel) and accretion (middle panel), and also their fraction (top panel) versus enclosed gas mass ($M_{\text{enc}}$) for the metallicities tested. The values were calculated just before the first sink particle was formed.

Fig. 8.—Timescales for fragmentation and accretion for different metallicities. "$t_{\text{frag}}(\langle N/\left(dN/dt\right)\rangle)$" indicates the average for the number of sink particles ($N$) divided by the time variation of that number, or the sink particle formation rate. "$t_{\text{acc}}(\langle M/\left(dM/dt\right)\rangle)$" is the average accretion time, which is calculated by dividing the total mass in sink particles divided by the mass accretion rate.
Evolution of temperatures in prestellar cloud cores with metallicities

Let us summarize here the formation processes of H

\[ T = \frac{1}{\gamma} \left( \frac{P}{\rho} \right) \]

where \( T \) is the temperature, \( P \) is the pressure, \( \rho \) is the density, and \( \gamma \) is the specific heat ratio.

EOS as function of metallicity

indicate the constant Jeans masses. For those above 10\(^\odot\)⊙, which is calculated by one-zone models. The dashed lines

Figure 1.

play a crucial role in the thermal evolution. The evolution of H

of specific heat at the center, one-zone models are presented. In Figure

compared with Figure 2 of O05, where similar plots for the

overall evolution is quite similar to that calculated by the one-

cases with metals. Below 10\(^{-3}\)\(\odot\), while fragmentation is strongly prohibited for 10\(^{-1}\)\(\odot\)⊙,

et al.

evolution. For example, the clouds easily fragment as long as

effective ratio of specific heat is an important index to examine

the core evolution. There are, however, small disagreements,

in particular, at high densities and for low-metallicity cases.

The metal-free one except for a slight offset at highest densities

are so small that the temperature evolution is almost identical to

later in Section

10

−3. PRESTELLAR COLLAPSE

\[ \tau \propto M_{\text{core}} \]

\[ \tau = 10^{-4} M_{\odot} \]

\[ \tau = 10^{-2} M_{\odot} \]

\[ \tau = 1 M_{\odot} \]

\[ \tau = 10^2 M_{\odot} \]

\[ \tau = 10^4 M_{\odot} \]

\[ \tau = 10^6 M_{\odot} \]

\[ \text{number density log (n_H (cm^{-3}))} \]

\[ \text{temperature T(K)} \]

(Omukai et al. 2005, 2010)
• slope of EOS in the density range $5 \times 10^3 \leq n \leq 16 \times 10^3$ cm$^{-3}$ is $\gamma \approx 1.06$.

• with non-zero angular momentum, disk forms.

• disk is unstable against fragmentation at high density

(Omukai et al. 2005, 2010)
Figure 1: Density evolution in a 120 AU region around the first protostar, showing the build-up of the protostellar disk and its eventual fragmentation. We also see ‘wakes’ in the low-density regions, produced by the previous passage of the spiral arms.
important disk parameters

Figure 2: Radial profiles of the disk's physical properties, centered on the first protostellar core to form. The quantities are mass-weighted and taken from a slice through the midplane of the disk. In the lower right-hand plot we show the radial distribution of the disk's Toomre parameter, $Q = \frac{c_s \kappa}{\pi G \Sigma}$, where $c_s$ is the sound speed and $\kappa$ is the epicyclic frequency. Because our disk is Keplerian, we adopted the standard simplification, and replaced $\kappa$ with the orbital frequency.

The molecular fraction is defined as the number density of hydrogen molecules ($n_{H_2}$), divided by the number density of hydrogen nuclei ($n$), such that fully molecular gas has a value of 0.5 (Clark et al. 2011b, Science, 331, 1040).

Toomre $Q$: $Q = \frac{c_s \kappa}{\pi G \Sigma}$

instability for $Q < 1$

(Clark et al. 2011b, Science, 331, 1040)
Most recent calculations:

fully sink-less simulations, following the disk build-up over ~10 years
(resolving the protostars - first cores - down to $10^5$ km ~ 0.01 $R_\odot$)

expected mass spectrum

![Graph showing mass spectrum]

we see “flat” mass spectrum

also talk by Athena Stacy
expected mass spectrum

• expected IMF is flat and covers a wide range of masses

• implications
  - because slope > -2, most mass is in massive objects as predicted by most previous calculations
  - most high-mass Pop III stars should be in binary systems
    --> source of high-redshift gamma-ray bursts
  - because of ejection, some low-mass objects (< 0.8 M⊙) might have survived until today and could potentially be found in the Milky Way

• consistent with abundance patterns found in second generation stars
The metallicities of extremely metal-poor stars in the halo are consistent with the yields of core-collapse supernovae, i.e. progenitor stars with 20 - 40 M⊙

(e.g. Tominaga et al. 2007, Izutani et al. 2009, Joggerst et al. 2009, 2010)
primordial star formation

- just like in present-day SF, we expect
  - turbulence
  - thermodynamics (i.e. heating vs. cooling)
  - feedback
  - magnetic fields
to influence first star formation.

- masses of first stars still uncertain, but we expect a wide mass range with typical masses of several 10s of $M_\odot$

- disks unstable: first stars in binaries or part of small clusters

- current frontier: include feedback and magnetic fields and possibly dark matter annihilation...
reducing fragmentation

• from present-day star formation theory we know, that
  can influence the fragmentation behavior.

• in the context of Pop III
  - magnetic fields: Turk et al. 2012, but see also Bovino et al. 2013

• all these will reduce degree of fragmentation
  (but not by much, see Rowan Smith et al. 2011, 2012, at least for accretion heating)

• DM annihilation might become important for disk dynamics and fragmentation
  (Ripamonti et al. 2011, Stacy et al. 2012b, Rowan Smith et al. 2012)
Star formation is intrinsically a multi-scale and multi-physics problem. Many different processes need to be considered simultaneously.
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- stars form from the competition of processes (such as pressure, CR pressure,
- thermodynamic properties in the star formation process
- detailed studies require the physical processes
- star formation is poorly understood
- primordial star formation
... people in the star formation group at Heidelberg University:

Christian Baczynski, Erik Bertram, Frank Bigiel, Andre Bubel, Diane Cormier, Volker Gaibler, Simon Glover, Dimitrious Gouliermis, Tilman Hartwig, Juan Ibanez, Christoph Klein, Lukas Konstandin, Mei Sasaki, Jennifer Schober, Rahul Shetty, Rowan Smith, László Szűcs

... former group members:

Robi Banerjee, Ingo Berentzen, Paul Clark, Christoph Federrath, Philipp Girichidis, Thomas Greif, Milica Micic, Thomas Peters, Dominik Schleicher, Stefan Schmeja, Sharanya Sur, ...

... many collaborators abroad!
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thanks