Molecular Cloud Fragmentation and Star Formation

Ralf Klessen

Zentrum für Astronomie der Universität Heidelberg
Institut für Theoretische Astrophysik
thanks to ...

many thanks to the members of the star formation group at the Institute for Theoretical Astrophysics at the Center for Astronomy of Heidelberg University

- Robi Banerjee
- Paul Clark
- Christoph Federrath
- Simon Glover
- Thomas Greif
- Susanne Horn
- Stefan Schmeja
- Thomas Peters
- Dominik Schleicher
- and many guests
formation of molecular clouds
  - on galactic scales
  - locally, in convergent flows
  - do molecular clouds lose memory of initial conditions?
fragmentation of molecular clouds
  - interplay between gravity and turbulence
star formation
  - initial mass function (models & caveats)
  - some examples
what’s next?
gravoturbulent star formation

idea:

Star formation is controlled by interplay between gravity and supersonic turbulence!

dual role of turbulence:

- stability on large scales
- initiating collapse on small scales

(e.g., Larson, 2003, Rep. Prog. Phys., 66, 1651; or Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125)
gravoturbulent star formation

idea:

Star formation is controlled by interplay between gravity and supersonic turbulence!

validity:

This hold on all scales and applies to build-up of stars and star clusters within molecular clouds as well as to the formation of molecular clouds in galactic disk.

(e.g., Larson, 2003, Rep. Prog. Phys, 66, 1651; or Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125)
gravoturbulent star formation

- interstellar gas is highly inhomogeneous
  - gravitational instability
  - thermal instability
  - turbulent compression (in shocks $\delta \rho/\rho \propto M^2$; in atomic gas: $M \approx 1...3$)

- cold molecular clouds can form rapidly in high-density regions at stagnation points of convergent large-scale flows
  - chemical phase transition: atomic $\rightarrow$ molecular
  - process is modulated by large-scale dynamics in the galaxy

- inside cold clouds: turbulence is highly supersonic ($M \approx 1...20$)
  - turbulence creates large density contrast, gravity selects for collapse

- turbulent cascade: local compression within a cloud provokes collapse
  - formation of individual stars and star clusters

(e.g. Mac Low & Klessen, 2004, Rev. Mod. Phys., 76, 125-194)
predictions

star formation on galactic scales
- global correlations: Schmidt-law
- efficiencies, rates, timescales, and long-term evolution:
  starburst vs. low surface density gal.
- triggers of star formation on global scales
- formation of dense cold molecular clouds
  properties of these clouds (structure, turbulence, etc.)

star cluster formation within clouds
- SF efficiency and timescale
- properties of young star clusters (structure, kinematics)
- stellar mass function – IMF
- multiplicity
- effects of stellar feedback (jets, outflows, radiation, winds, ...)

Ralf Klessen: Ringberg 29.07.2008
molecular cloud formation

- star formation on galactic scales
  \rightarrow missing link so far:
  \textit{formation of molecular clouds}

- questions
  
  - where and when do molecular clouds form?
  
  - what are their properties?
  
  - how does that correlation to star formation?
  
  - global correlations? \rightarrow Schmidt law
Thesis:
Molecular clouds form at *stagnation points* of large-scale convergent flows, mostly triggered by global (or external) perturbations.

(Deul & van der Hulst 1987, Blitz et al. 2004)
correlation with large-scale perturbations

density/temperature fluctuations in warm atomar ISM are caused by thermal/gravitational instability and/or supersonic turbulence

some fluctuations are dense enough to form $H_2$ within “reasonable time”

$\rightarrow$ molecular cloud

(Glover & Mac Low 2007a,b)

external perturbuations (i.e. potential changes) increase likelihood

(e.g. talk by Clare Dobbs)

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star formation on *global* scales

probability distribution function of the density ($\rho$-pdf)

varies rms Mach numbers:

- $M_1 > M_2$
- $M_3 > M_4 > 0$

mass weighted $\rho$-pdf, each shifted by $\Delta \log N = 1$

(from Klessen, 2001; also Gazol et al. 2005, Mac Low et al. 2005)
star formation on global scales

H₂ formation rate:

\[ \tau_{H_2} \approx \frac{1.5 \text{ Gyr}}{n_{H}/1 \text{ cm}^{-3}} \]

for \( n_{H} \geq 100 \text{ cm}^{-3} \), H₂ forms within 10 Myr, this is about the lifetime of typical MC’s.

in turbulent gas, the H₂ fraction can become very high on short timescale

(for models with coupling between cloud dynamics and time-dependent chemistry, see Glover & Mac Low 2007a,b)

mass weighted \( \rho \)-pdf, each shifted by \( \Delta \log N = 1 \)

(rate from Hollenback, Werner, & Salpeter 1971)
modeling galactic SF

SPH calculations of self-gravitating disks of stars and (isothermal) gas in dark-matter potential, sink particles measure local collapse --> star formation
We find a correlation between star formation rate and gas surface density:

$$\Sigma_{SFR} \propto \Sigma_{gas}^{1.5}$$

**global Schmidt law**
in both cases: \[ \Sigma_{\text{SFR}} \propto \Sigma_{\text{gas}}^{1.5} \]

(from Kennicutt 1998)
molecular cloud formation

(from Dobbs, Glover, Clark, Klessen 2008)
molecular cloud formation

molecular gas fraction as function of time

- $\Sigma=4$ $M_{\odot}$ pc$^{-2}$
- $\Sigma=10$ $M_{\odot}$ pc$^{-2}$
- $\Sigma=20$ $M_{\odot}$ pc$^{-2}$

molecular gas fraction as function of density

(Dobbs et al. 2008)
molecular cloud formation

molecular gas fraction of fluid element as function of time

molecular gas fraction as function of density

(Dobbs et al. 2008)
observed timescales

Tamburro et al. (2008)

Fig. 1.— NGC 5194: the 24 μm band image is plotted in color scale; the H1 emission map is overlayed with green contours.
observed timescales

Tamburro et al. (2008)

Fig. 5.— Histogram of the time scales $t_{\text{HI} \rightarrow 24\,\mu\text{m}}$ derived from the fits in Figure 4 and listed in Table 2 for the 14 sample galaxies listed in Table 1. The timescales range between 1 and 4 Myr for almost all galaxies.
consistent models of ISM dynamics require to go beyond the simple models!

- magnetohydrodynamics (account for large-scale dynamics + turbulence)
- time-dependent chemistry (reduced network, focus on few dominant species, e.g. H$_2$)
- radiation (currently simple assumptions)

H$_2$ forms rapidly in shocks / transient density fluctuations / H$_2$ gets destroyed slowly in low density regions / result: turbulence greatly enhances H$_2$-formation rate

(Glover & Mac Low 2007ab:)

(ISM: transition H I to H$_2$)
Reduced chemical network

Table 1. The set of chemical reactions that make up our model of non-equilibrium hydrogen chemistry.

<table>
<thead>
<tr>
<th>Reaction</th>
<th>Reference</th>
</tr>
</thead>
<tbody>
<tr>
<td>1. $, \text{H} + \text{H} + \text{grain} \rightarrow \text{H}_2 + \text{grain}$</td>
<td>Hollenbach &amp; McKee (1979)</td>
</tr>
<tr>
<td>2. $, \text{H}_2 + \text{H} \rightarrow 3\text{H}$</td>
<td>Mac Low &amp; Shull (1986) (low density), Lepp &amp; Shull (1983) (high density)</td>
</tr>
<tr>
<td>3. $, \text{H}_2 + \text{H}_2 \rightarrow 2\text{H} + \text{H}_2$</td>
<td>Martin, Krogh &amp; Mundy (1998) (low density), Shapiro &amp; Kang (1987) (high density)</td>
</tr>
<tr>
<td>4. $, \text{H}_2 + \gamma \rightarrow 2\text{H}$</td>
<td>See § 2.2.1</td>
</tr>
<tr>
<td>5. $, \text{H} + \text{e} \rightarrow \text{H}^+ + \text{e}$</td>
<td>Liszt (2003)</td>
</tr>
<tr>
<td>6. $, \text{H} + \text{e} \rightarrow \text{H}^+ + 2\text{e}$</td>
<td>Abel et al. (1997)</td>
</tr>
<tr>
<td>7. $, \text{H}^+ + \text{e} \rightarrow \text{H} + \gamma$</td>
<td>Ferland et al. (1992)</td>
</tr>
<tr>
<td>8. $, \text{H}^+ + \text{e} + \text{grain} \rightarrow \text{H}^+ + \text{grain}$</td>
<td>Weingartner &amp; Draine (2001)</td>
</tr>
</tbody>
</table>

Table 2. Processes included in our thermal model.

<table>
<thead>
<tr>
<th>Process</th>
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<tbody>
<tr>
<td></td>
<td>Collisional rates ($\text{H}_2$) – Flower &amp; Lunney (1977)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates ($\text{H}, T &lt; 2000 \text{ K}$) – Hollenbach &amp; McKee (1989)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates ($\text{H}, T &gt; 2000 \text{ K}$) – Keenan et al. (1986)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates ($\text{e}^-$) – Wilson &amp; Bell (2002)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates ($\text{H}, \text{H}_2$) – Flower, priv. comm.</td>
</tr>
<tr>
<td></td>
<td>Collisional rates ($\text{e}^-$) – Bell, Berrington &amp; Thomas (1958)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates ($\text{H}^+$) – Pequignot (1990, 1996)</td>
</tr>
<tr>
<td></td>
<td>Collisional rates ($\text{H}_2$, $\text{H}_2^+$) – Flower, priv. comm.</td>
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<tr>
<td></td>
<td>Collisional rates ($\text{e}^-$) – Dufour &amp; Kingson (1991)</td>
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<td>Collisional rates ($\text{H}$) – Houff (1990)</td>
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<td></td>
<td>Collisional rates ($\text{e}^-$) – Dufour &amp; Kingson (1991)</td>
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<td></td>
<td>Gas-grain energy transfer$^1$ – Le Bourlot, Pineau des Forêts &amp; Flower (1999)</td>
</tr>
<tr>
<td></td>
<td>Holmberg &amp; McKee (1989)</td>
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<tr>
<td></td>
<td>Recombination on grains – Wilson et al. (2003)</td>
</tr>
<tr>
<td></td>
<td>Atomic data – Sutherland &amp; Dalgarno (1983)</td>
</tr>
<tr>
<td></td>
<td>H$^+$ collisional ionization – Abel et al. (1997)</td>
</tr>
<tr>
<td></td>
<td>$\text{H}_2$ collisional dissociation – See Table 1</td>
</tr>
</tbody>
</table>

Heating:

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<tr>
<td>Photoelectric effect</td>
<td>Basco &amp; Tielens (1994); Wolfire et al. (2003)</td>
</tr>
<tr>
<td>$\text{H}_2$ photodissociation</td>
<td>Black &amp; Dalgarno (1977)</td>
</tr>
<tr>
<td>UV pumping of $\text{H}_2$</td>
<td>Burton, Hollenbach &amp; Tielens (1990)</td>
</tr>
<tr>
<td>$\text{H}_2$ formation on dust grains</td>
<td>Hollenbach &amp; McKee (1989)</td>
</tr>
<tr>
<td>Cosmic ray ionization</td>
<td>Goldsmith &amp; Langer (1978)</td>
</tr>
</tbody>
</table>

here: $\text{e}^-$, $\text{H}^+$, $\text{H}$, $\text{H}_2$

in primordial gas we do:

$\text{e}^-$, $\text{H}^+$, $\text{H}$, $\text{H}_2$, $\text{H}_2^+$, $\text{H}_2$, $\text{C}$, $\text{C}^+$, $\text{O}$, $\text{O}^+$

(Glover & Mac Low 2007ab)
$L = 40 \text{ pc}$, $n_0 = 100 \text{ cm}^{-3}$, $B_0 = 5.85 \text{ mG}$, $v_{\text{rms}} = 0.0$

(Glover & Mac Low 2007a)
L = 20 pc, \( B_0 = 5.85 \, \mu G \), \( v_{\text{rms}} = 10 \, \text{km/s} \)

(Glover & Mac Low 2007a)
from atomic gas to molecular clouds

thesis: **cold molecular clouds can form rapidly in high-density regions at stagnation points of convergent large-scale flows**

- chemical phase transition: atomic $\rightarrow$ molecular
- process is modulated by large-scale dynamics in galaxy

questions

- are molecular clouds truly “multi-phase” media?
- turbulence? dynamical & morphological properties?
- what is relation to initial & environmental conditions?
- magnetic field structure? (see special discussion session)
convergent flows: set-up

- convergent flow studies
  - atomic flows collide
  - cooling curve (soon chemistry)
  - gravity
  - magnetic fields
  - numerics: AMR, BGK, SPH

from Vazquez-Semadeni et al. (2007) see studies by Banerjee et al., Heitsch et al., Hennebelle et al., Vazquez-Semadeni et al.
convergent flows: set-up

- convergent flow studies
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from Vazquez-Semadeni et al. (2007)

see studies by Banerjee et al., Heitsch et al., Hennebelle et al., Vazquez-Semadeni et al.
MC formation in convergent flows

thermal instability + gravity creates complex molecular cloud structure:

from Banerjee et al. (2008)
(see also studies by Hennebelle et al. and Vazquez-Semadeni et al. as well as talk by Fabian Heitsch)
some results: density & B-field

Fig. 2. Top left panel: column density. Top right panel: Magnetic intensity and its xy-components (indicated as arrows) in the $z = 0$ plane. Bottom left panel: density and velocity fields in the $z = 0$ plane. Bottom right panel: temperature in the $z = 0$ plane.

from Hennebelle et al. (2008) initially $B = 5 \mu G$
some results: density & B-field

from Hennebelle et al. (2008)

initially $B = 5 \mu G$

$B \sim n^{1/2}$
some results: growth of cores

**Figure 2.** Shows the time evolution of a typical clump which initially develops out of the thermally unstable WNM in shock layers of turbulent flows. A small cold condensate grows by outward propagation of its boundary layer. Coalescence and merging with nearby clumps further increases the size and mass of these clumps. The global gravitational potential of the proto-cloud enhances the merging probability with time. The images show 2D slices of the density (logarithmic colour scale) and the gas velocity (indicated as arrows) in the plane perpendicular to the large scale flows.

**two phases of core growth:**
(1) by *outward propagation of boundary layer* → Jeans sub-critical phase
(2) *core mergers* → super-Jeans → gravitational collapse & star formation
example: *Pipe nebula* ???

from Banerjee et al. (2008)
relation between flow and magnetic field:
mass flow mostly along field lines

from Banerjee et al. (2008)
relation between flow and magnetic field:

mass flow mostly along field lines

from Banerjee et al. (2008)
some results: growth of cores

Figure 3. Shows the structure of one typical clump which forms in the thermally unstable WNM gas. The images show 2D slices of the temperature (left, log scale), thermal pressure (middle, linear scale), and magnetic field strength (right, linear scale). The arrows in the temperature and pressure plots indicate the velocity field and, in the right panel, the magnetic stream lines. The cold ($T \sim 30 - 50 K$), dense ($n \sim 2 - 5 \times 10^3 \text{ cm}^{-3}$) molecular clump is embedded in the warm atomic gas ($T \sim 5 \times 10^3 K$) and has a well defined boundary. Due to the thermal properties of the ISM (see Fig. 2 of Vázquez-Semadeni et al. 2007, for the equilibrium pressure), such clumps are almost in pressure equilibrium with their surrounding. The overdense clumps exert a gravitational force on the low density environment where gas continues to stream into the clump predominately anti-parallel to the magnetic flux lines (see also Fig. 2).

**some properties of cores:**

1. cores are in approximate pressure equilibrium with surrounding
2. accretion / mass flow mostly along magnetic field lines
3. core densities $n \sim 2 - 5 \times 10^3 \text{ cm}^{-3}$, core temperature $T \sim 30 - 50 K$

from Banerjee et al. (2008)
fragmentation of molecular clouds and relation to stellar birth

some questions

- how does the turbulence generated by cloud formation influence cloud fragmentation?
- how important is turbulence from internal feedback? (is that consistent with observations?)
- interplay between gravity and turbulence? → role of turbulence for star formation
star-forming filaments in the *Taurus* cloud
(from Alyssa Goodman)

Structure and dynamics of young star clusters is coupled to *structure of molecular cloud*
what drives turbulence?

- turbulence characteristics
  - molecular cloud turbulence seems to be dominated by large-scale models
  - consistent with external driving
  - convergent flows?
    → the same process that creates the cloud supplies internal turbulence ...
  - alternative mechanisms:
    - gravity (spiral shocks), supernovae, HII regions?
    - internal sources: jets, outflows?

Polaris flare (from Ossenkopf & Mac Low 2002)
what drives turbulence?

- turbulence characteristics
  - molecular cloud turbulence seems to be dominated by large-scale models
  - consistent with external driving
  - **convergent flows?**
    - the same process that creates the cloud supplies internal turbulence ..
    - caused by
      - gravity (spiral shocks), supernovae, HII regions?
  - alternative mechanisms:
    - internal sources: jets, outflows?

Polaris flare (from Ossenkopf & Mac Low 2002)
what drives turbulence?

- some words on internal sources
  - molecular cloud turbulence seems to be dominated by large-scale models
  - jets / outflow can only work after onset of star formation
    → what about turbulence in non-star forming parts of clouds, or during initial phases?

- there is debate on effectiveness of internal sources for driving supersonic turbulence
  (Li & Nakamura vs. Banerjee, Klessen, Fendt)

(Nakamura & Li 2007) (Banerjee, Klessen, Fendt 2008)
individual jets cannot drive supersonic turbulence in a space-filling way $\rightarrow$ need additional physics

Banerjee, Klessen, & Fendt (2008)
jets from cluster with self-gravity with AMR code FLASH

Banerjee et al. (very preliminary study)
turbulence
Turbulent cascade

Kolmogorov (1941) theory incompressible turbulence

Inertial range: scale-free behavior of turbulence

„Size“ of inertial range:
\[ \frac{L}{\eta_K} \approx \text{Re}^{3/4} \]

Energy input scale

Energy dissipation scale

Logarithmic scales:
- Log $E$ vs. Log $k$
Shock-dominated turbulence

Turbulent cascade

\[ \log E \quad \log k \]

Energy input scale \( L^{-1} \)

Energy dissipation scale \( \eta_K^{-1} \)

\[ k^{-2} \]

\[ \eta_K^{-1} \]

\[ L^{-1} \]

\[ \frac{L}{\eta_K} \approx Re^{3/4} \]

Inertial range: scale-free behavior of turbulence

"Size" of inertial range:
Turbulent cascade in ISM

- Energy source & scale: NOT known (supernovae, winds, spiral density waves?)
  - $\sigma_{\text{rms}} \ll 1 \text{ km/s}$
  - $M_{\text{rms}} \leq 1$
  - $L \approx 0.1 \text{ pc}$

- Dissipation scale not known (ambipolar diffusion, molecular diffusion?)

Diagram:
- Molecular clouds: $\sigma_{\text{rms}} \approx \text{several km/s}$, $M_{\text{rms}} > 10$, $L > 10 \text{ pc}$
- Massive cloud cores: $\sigma_{\text{rms}} \approx \text{few km/s}$, $M_{\text{rms}} \approx 5$, $L \approx 1 \text{ pc}$
- Dense protostellar cores: $\sigma_{\text{rms}} \ll 1 \text{ km/s}$, $M_{\text{rms}} \leq 1$, $L \approx 0.1 \text{ pc}$
Density structure of MC’s

Density structure of MC’s

molecular clouds are highly inhomogeneous

stars form in the densest and coldest parts of the cloud

ρ-Ophiuchus cloud seen in dust emission

let’s focus on a cloud core like this one

(Motte, André, & Neri 1998)

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Evolution of cloud cores

How does this core evolve?
Does it form one single massive star or cluster with mass distribution?

- Turbulent cascade “goes through” cloud core
  -- NO scale separation possible
  -- NO effective sound speed

- Turbulence is supersonic!
  -- produces strong density contrasts:
    \[ \frac{\delta \rho}{\rho} \approx M^2 \]
  -- with typical \( M \approx 10 \) -- \( \frac{\delta \rho}{\rho} \approx 100 \! \)!

- many of the shock-generated fluctuations are Jeans unstable and go into collapse

-- expectation: core breaks up and forms a cluster of stars
indeed ρ-Oph B1/2 contains several cores ("starless" cores are denoted by x, cores with embedded protostars by ★)

(Motte, André, & Neri 1998)
Formation and evolution of cores

- protostellar cloud cores form at the **stagnation points** of **convergent turbulent flows**
  - if \( M > M_{\text{Jeans}} \propto \rho^{-1/2} T^{3/2} \): collapse and star formation
  - if \( M < M_{\text{Jeans}} \propto \rho^{-1/2} T^{3/2} \): reexpansion after external compression fades away

- typical timescales: \( t \approx 10^4 \ldots 10^5 \text{ yr} \)

- because **turbulent** ambipolar diffusion time is **short**, this time estimate still holds for the presence of magnetic fields, in **magnetically critical cores**

  (e.g. Fatuzzo & Adams 2002, Heitsch et al. 2004)
Formation and evolution of cores

- protostellar cloud cores form at stagnation point in convergent turbulent flows

- if $M > M_{\text{crit}} \propto \rho^{-1/2} T^{3/2}$: collapse & star formation

- if $M < M_{\text{crit}} \propto \rho^{-1/2} T^{3/2}$: reexpansion after end of external compression

  (e.g. Vazquez-Semadeni et al 2005)

- typical timescale: $t \approx 10^4 \ldots 10^5$ yr
Formation and evolution of cores

What happens to distribution of cloud cores?

Two extreme cases:

1. Turbulence dominates energy budget:
   \[ \alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} > 1 \]
   ---> individual cores do not interact
   ---> collapse of individual cores dominates stellar mass growth
   ---> loose cluster of low-mass stars

2. Turbulence decays, i.e. gravity dominates: \[ \alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} < 1 \]
   ---> global contraction
   ---> cores do interact while collapsing
   ---> competition influences mass growth
   ---> dense cluster with high-mass stars
turbulence creates a hierarchy of clumps
as turbulence decays locally, contraction sets in
as turbulence decays locally, contraction sets in
while region contracts, individual clumps collapse to form stars
while region contracts, individual clumps collapse to form stars
individual clumps collapse to form stars
individual clumps collapse to form stars
\[ \alpha = \frac{E_{\text{kin}}}{|E_{\text{pot}}|} < 1 \]

In *dense clusters*, clumps may merge while collapsing.

--> then contain multiple protostars
in *dense clusters*, clumps may merge while collapsing
--> then contain multiple protostars
in *dense clusters*, clumps may merge while collapsing
--> 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
result: *star cluster* with H\text{I} region
initial mass function
initial mass function

what is the relation between molecular cloud fragmentation and the distribution of stars?

important quantity: **IMF**

BUT: “everyone” gets the right IMF → better look for secondary indicators

- **stellar multiplicity** (see focus group lead by H. Zinnecker)
- protostellar **spin** (including disk)
- **spatial distribution** + **kinematics** in young clusters
- **magnetic field strength** and **orientation**
  (see focus group lead by R. Crutcher)
distribution of stellar masses depends on

- turbulent initial conditions
  --> mass spectrum of prestellar cloud cores

- collapse and interaction of prestellar cores
  --> competitive 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

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
IMF

distribution of stellar masses depends on

- turbulent initial conditions
  --> mass spectrum of prestellar cloud cores ???

- collapse and interaction of prestellar cores
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(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
compressive vs. rotational driving

- statistical characteristics of turbulence depend strongly on “type“ of driving
- example: dilatational vs. solenoidal driving
- question: what drives ISM turbulence on different scales?
Federrath, Klessen, Schmidt (2008a)
dilatational vs. solenoidal

- density pdf depends on “dimensionality” of driving
  - relation between width of pdf and Mach number
    \[ \frac{\sigma_\rho}{\rho_0} = b \mathcal{M} \]
    - with \( b \) depending on \( \zeta \) via
      \[
      b = 1 + \left[ \frac{1}{D} - 1 \right] \zeta = \begin{cases} 
      1 - \frac{2}{3} \zeta, & \text{for } D = 3 \\
      1 - \frac{1}{2} \zeta, & \text{for } D = 2 \\
      1, & \text{for } D = 1 
      \end{cases}
      \]
    - with \( \zeta \) being the ratio of dilatational vs. solenoidal modes:
      \[
      \mathcal{P}^\zeta_{ij} = \zeta \mathcal{P}_{ij}^\perp + (1 - \zeta) \mathcal{P}_i^\parallel = \zeta \delta_{ij} + (1 - 2\zeta) \frac{k_i k_j}{|k|^2}
      \]

Federrath, Klessen, Schmidt (2008a)
dilatational vs. solenoidal

- density pdf depends on “dimensionality” of driving
  → is that a problem for the Krumholz & McKee model of the SF efficiency?

- density pdf of compressive driving is **NOT log-normal**
  → is that a problem for the Padoan & Nordlund IMF model?

- most “physical” sources should be **compressive** (convergent flows from spiral shocks or SN)

good fit needs 3rd and 4th moment of distribution!

Federrath, Klessen, Schmidt (2008b)
density power spectrum differs between dilatational and solenoidal driving!

\[ \Rightarrow \text{dilatational driving leads to break at sonic scale!} \]

can we use that to determine driving sources from observations?

compensated density spectrum \( kS(k) \) shows clear break at sonic scale. below that shock compression no longer is important in shaping the power spectrum ...
IMF

distribution of stellar masses depends on

- turbulent initial conditions
  --> mass spectrum of prestellar cloud cores

- **collapse and interaction of prestellar cores**
  --> competitive 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

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
example: model of Orion cloud

(Spitzer: Megeath et al.)

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

Bonnell & Clark 2008
Dynamics of nascent star cluster

in dense clusters protostellar interaction may be come important!

Trajectories of protostars in a nascent dense cluster created by gravoturbulent fragmentation
Mass accretion rates vary with time and are strongly influenced by the cluster environment.

distribution of stellar masses depends on

- turbulent initial conditions
  --> mass spectrum of prestellar cloud cores

- collapse and interaction of prestellar cores
  --> competitive 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

(e.g. Larson 2003, Prog. Rep. Phys.; Mac Low & Klessen, 2004, Rev. Mod. Phys, 76, 125 - 194)
dependency on EOS

- 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

dependency on EOS

for $\gamma<1$ fragmentation is enhanced $\rightarrow$ cluster of low-mass stars
for $\gamma>1$ it is suppressed $\rightarrow$ formation of 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) \( M_{\text{jeans}} \propto \gamma^{3/2} \rho^{(3\gamma-4)/2} \)

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

- \( \gamma > 1 \): \( \Rightarrow \text{small} \) density excursion for given pressure
  \( \Rightarrow \langle M_{\text{jeans}} \rangle \) is large
  \( \Rightarrow \) only few and massive clumps exceed \( M_{\text{jeans}} \)
EOS for solar neighborhood

below $10^{-18}$ gcm$^{-3}$: $\rho \uparrow \rightarrow T \downarrow$

above $10^{-18}$ gcm$^{-3}$: $\rho \uparrow \rightarrow T \uparrow$

$Larson$ 1985, Larson 2005

$P \propto \rho^{\gamma}$

$P \propto \rho T$

$\rightarrow \gamma = 1 + \frac{d\ln T}{d\ln \rho}$

$\gamma = 0.7$

$\gamma = 1.1$
IMF from simple piece-wise polytropic EOS

\( \gamma_1 = 0.7 \)

\( \gamma_2 = 1.1 \)

\[ T \sim \rho^{\gamma-1} \]

(Jappsen et al. 2005)
IMF from simple piece-wise polytropic EOS

critical density $\uparrow$ median mass $\downarrow$

(Jappsen et al. 2005)
IMF in nearby molecular clouds

with $\rho_{\text{crit}} \approx 2.5 \times 10^5 \text{ cm}^{-3}$ at SFE $\approx 50\%$

Isothermal EOS has deficits of very low-mass objects

--> need "better" EOS!
Plausibility argument for shape

Supersonic turbulence is scale free process

→ POWER LAW BEHAVIOR

But also: turbulence and fragmentation are highly stochastic processes → central limit theorem

→ GAUSSIAN DISTRIBUTION
metallicity dependence
EOS as function of metallicity

(Omukai et al. 2005)
EOS as function of metallicity

(Omukai et al. 2005)
EOS as function of metallicity

(Omukai et al. 2005)
present-day star formation

present-day star formation

- **fragmentation behavior** depends on **EOS**

- **"kinck" in EOS** introduces **characteristic mass**

- **IMF** depends on **scale (and strength) of turbulence**

- **characteristic mass** is (relatively) **insensitive to environmental parameters** → **universal IMF in local universe**
dependence on Z at low density

\((\text{Omukai et al. 2005})\)
dependence on Z at low density

- at densities below \( n \approx 10^2 \text{ cm}^{-3} \), \( \text{H}_2 \) cooling dominates the behavior. (Jappsen et al. 2007)

- fragmentation depends on initial conditions. then

  - example: solid-body rotating top-hat initial conditions with dark matter fluctuations (a la Bromm et al. 1999) fragment no matter what metallicity you take (in regime \( n \leq 10^6 \text{ cm}^{-3} \)) → because unstable disk builds up

  (Jappsen et al. 2008a)
dependence on Z at low density

\[ Z = 0 \]

rotating top-hat with dark matter fluctuations fragments, no matter what

\[ Z = -4 \]

\[ Z = -3 \]

\[ Z = -2 \]

\[ Z = -1 \]

(Jappsen et al. 2008a, see also Clark et al. 2008)
dependence on $Z$ at low density

- fragmentation depends on *initial conditions*

  then

- example: *centrally concentrated halo* does not fragment up to densities of $n \approx 10^6 \text{ cm}^{-3}$ up to metallicities $Z \approx -1$

  (Jappsen et al. 2008b)
slope of EOS in the density range $5 \text{ cm}^{-3} \leq n \leq 16 \text{ cm}^{-3}$ is $\gamma \approx 1.06$.

- with non-zero angular momentum, disk forms.
- disk is unstable against fragmentation at high density
transition: Pop III to Pop II.5

(Omukai et al. 2005)
dust induced fragmentation at $Z=10^{-5}$

$\tau = t_{\text{SF}} - 67 \text{ yr}$
$\tau = t_{\text{SF}} - 20 \text{ yr}$
$\tau = t_{\text{SF}}$

$\tau = t_{\text{SF}} + 53 \text{ yr}$
$\tau = t_{\text{SF}} + 233 \text{ yr}$
$\tau = t_{\text{SF}} + 420 \text{ yr}$

(Clark et al. 2007)
Ralf Klessen: Ringberg 29.07.2008
dense cluster of low-mass protostars builds up:

- mass spectrum peaks below $1 \, M_{\odot}$
- cluster VERY dense
  $n_{\text{stars}} = 2.5 \times 10^9 \, \text{pc}^{-3}$
- fragmentation at density
  $n_{\text{gas}} = 10^{12} - 10^{13} \, \text{cm}^{-3}$

(Clark et al. 2007)
cluster build-up

\[ \text{Fig. 3.} \quad \text{We illustrate the onset of the fragmentation process in the high resolution } Z = 10^{-5} \text{ Z}_\odot \text{ simulation. The graphs show the densities of the particles, plotted as a function of their x-position. Note that for each plot, the particle data has been centered on the region of interest. We show here results at three different output times, ranging from the time that the first star forms (} t_{SF} \text{) to 221 years afterwards. The densities lying between the two horizontal dashed lines denote the range over which dust cooling lowers the gas temperature.} \]

(Clark et al. 2007)
cluster build-up

\( \gamma > 1 \) (heating)

\( \gamma < 1 \) (cooling)

(Clark et al. 2007)
gas properties at time when first star forms

(Clark et al. 2007)
dense cluster of low-mass protostars builds up:

- mass spectrum peaks below $1\,\text{M}_{\text{sun}}$
- cluster VERY dense $n_{\text{stars}} = 2.5 \times 10^9\,\text{pc}^{-3}$
- fragmentation at density $n_{\text{gas}} = 10^{12} - 10^{13}\,\text{cm}^{-3}$

(Clark et al. 2007)
Summary I

- interstellar gas is highly *inhomogeneous*
  - *thermal instability*
  - *gravitational instability*
  - *turbulent compression* (in shocks $\delta \rho/\rho \propto M^2$; in atomic gas: $M \approx 1...3$)

- cold *molecular clouds* can form rapidly in high-density regions at *stagnation points* of convergent large-scale flows
  - chemical *phase transition*: atomic $\rightarrow$ molecular
  - process is *modulated* by large-scale dynamics in the galaxy

- inside *cold clouds*: turbulence is highly supersonic ($M \approx 1...20$)
  - *turbulence* creates density contrast, *gravity* selects for collapse
    - GRAVOTUBULENT FRAGMENTATION

- *turbulent cascade*: local compression within a cloud provokes collapse $\rightarrow$ formation of individual *stars* and *star clusters*

- *star cluster*: gravity dominates in large region ($\rightarrow$ competitive accretion)

(e.g. Mac Low & Klessen, 2004, Ballesteros-Paredes et al. 2006, McKee & Ostriker 2007)

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Summary II

- **thermodynamic response** (EOS) determines fragmentation behavior
  - characteristic stellar mass from fundamental atomic and molecular parameters
    --> explanation for quasi-universal IMF?

- **stellar feedback** is important
  - accretion heating may reduce degree of fragmentation
  - ionizing radiation will set efficiency of star formation

- **CAVEATS:**
  - star formation is *multi-scale, multi-physics* problem --> VERY difficult to model
  - in simulations: very small turbulent inertial range (Re < 1000)
  - can we use EOS to describe thermodynamics of gas, or do we need time-dependent chemical network and radiative transport?
  - stellar feedback requires (at least approximative) radiative transport, most numerical calculations so far have neglected that aspect

(e.g. Mac Low & Klessen, 2004, Ballesteros-Paredes et al. 2006, McKee & Ostriker 2007)
LEBEN

PLANETENSYSTEME

1. Moleküle
   - Physikalische Chemie
   - Atomphysik
   - Komplexität
   - Intergalaktische Materie

2. Planetenentstehung

3. Sternentstehung

4. Protostern

5. Staub + Gas

6. Kleine Sternmasse
   - Planetarische Nebel
   - Gas + Staub

7. Große Sternmasse
   - Supernovae
   - Weiße Zwerge
   - Neutronensterne
   - Schwarze Löcher

Ralf Klessen: Ringberg 29.07.2008
what do we need ... 
... to study ISM and star formation?

- **magneto-hydrodynamics**  
  (multi-phase, non-ideal MHD, turbulence)

- **chemistry** (gas + dust, heating + cooling)

- **radiation** (continuum + lines)

- **stellar dynamics**  
  (collisional: star clusters, collisionless: galaxies, DM)

- **stellar evolution**  
  (feedback: radiation, winds, SN)

+ **laboratory work**  
  (reaction rates, cross sections, dust coagulation properties, etc.)
what do we need?

- massive parallel codes
  - particle-based: SPH with improved algorithms (XSPH with turb. subgrid model, GPM, particle splitting, MHD-SPH?)
  - grid-based: AMR (FLASH, ENZO, RAMSES, Nirvana3, etc), subgrid-scale models (FEARLESS)
  - BGK methods

- magneto-hydrodynamics
  - (multi-phase, non-ideal MHD, turbulence)

- chemistry
  - (gas + dust, heating + cooling)

- radiation
  - (continuum + lines)

- stellar dynamics
  - (collisional: star clusters, collisionless: galaxies, DM)

- stellar evolution
  - (feedback: radiation, winds, SN)
what do we need?

- ever increasing chemical networks
- working reduced networks for time-dependent chemistry in combination with hydrodynamics
- improved data on reaction rates (laboratory + quantum mechanical calculations)

magneto-hydrodynamics
  (multi-phase, non-ideal MHD, turbulence)

chemistry (gas + dust, heating + cooling)

radiation (continuum + lines)

stellar dynamics
  (collisional: star clusters, collisionless: galaxies, DM)

stellar evolution
  (feedback: radiation, winds, SN)
what do we need?

- continuum vs. lines
- Monte Carlo, characteristics
- approximative methods
- combine with hydro

magneto-hydrodynamics
(multi-phase, non-ideal MHD, turbulence)

chemistry (gas + dust, heating + cooling)

radiation (continuum + lines)

stellar dynamics
(collisional: star clusters, collisionless: galaxies, DM)

stellar evolution
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what do we need?

- **magneto-hydrodynamics**
  (multi-phase, non-ideal MHD, turbulence)

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- **radiation** (continuum + lines)

- **stellar dynamics**
  (collisional: star clusters, collisionless: galaxies, DM)

- **stellar evolution**
  (feedback: radiation, winds, SN)

- statistics: number of stars (collisional: $10^6$, collisionless: $10^{10}$)
- transition from gas to stars
- binary orbits
- long-term integration
what do we need?

- magneto-hydrodynamics (multi-phase, non-ideal MHD, turbulence)
- chemistry (gas + dust, heating + cooling)
- radiation (continuum + lines)
- stellar dynamics (collisional: star clusters, collisionless: galaxies, DM)
- stellar evolution (feedback: radiation, winds, SN)

- very early phases (pre main sequence tracks)
- massive stars at late phases
- role of rotation
- primordial star formation
what do we need?

- **magneto-hydrodynamics**
  (multi-phase, non-ideal MHD, turbulence)

- **chemistry** (gas + dust, heating + cooling)

- **radiation** (continuum + lines)

- **stellar dynamics**
  (collisional: star clusters, collisionless: galaxies, DM)

- **stellar evolution**
  (feedback: radiation, winds, SN)

- **laboratory work**
  (reaction rates, cross sections, dust coagulation properties, etc.)

methods need to be combined!

Ralf Klessen: Ringberg 29.07.2008
Barnard 68: a well-studied isolated prestellar core

(Lada et al. 2003)
Barnard 68

\( \text{C}^{18}\text{O} \ (1-0) \)

\( \text{N}_2\text{H}^+ \ (1-0) \)

\( \text{CS} \ (3-2) \)

(Bergin et al.)

Ralf Klessen: Ringberg 29.07.2008
MHD model with proper heating and cooling terms (EOS)

chemical model

line radiative transfer

synthetic images of model cores

"Taste Testing Initiative → special discussion round lead by Alyssa Goodman"
Thanks!